WSRT Polarimetric Imaging of the warm ISM

Proefschrift

ter verkrijging van
de graad van Doctor aan de Universiteit Leiden,
op gezag van de Rector Magnificus Dr. D.D. Breimer,
hoogleraar in de faculteit der Wiskunde en
Natuurwetenschappen en die der Geneeskunde,
volgens besluit van het College voor Promoties
te verdedigen op donderdag 19 december 2002
te klokke 15.15 uur

doors

Marijke Haverkorn van Rijsewijk

geboren te Oegstgeest in 1974
Promotiecommissie

Promotores
Prof. dr. G. Miley
Prof. dr. A. G. de Bruyn (ASTRON, Dwingeloo; Kaptejin Instituut, Groningen)

Co-promotor
Dr. P. Katgert

Referent
Prof. dr. B. J. Rickett (Universiteit van Californië, San Diego)

Overige leden
Dr. R. Beck (Max-Planck-Instituut voor Radioastronomie, Bonn)
Prof. dr. H. J. Habing
Prof. dr. V. Icke
Prof. dr. R. T. Schilizzi (Sterrewacht Leiden; JIVE, Dwingeloo)
## Contents

1 Introduction .......................... 1
   1.1 The gaseous component of the ISM .................. 2
   1.2 The local ISM ............................ 4
   1.3 The Galactic magnetic field ....................... 5
   1.4 Cosmic rays and synchrotron radiation ............... 7
   1.5 ISM in external galaxies ....................... 8
   1.6 Polarization as a probe of structure in the warm ISM . 8
   1.7 Outline of the thesis ........................ 9
   1.8 Main results ............................... 11

2 WSRT Polarimetry .................... 15
   2.1 Observing with the WSRT ...................... 15
   2.2 Calibration of polarization data ................ 16
   2.3 Constructing maps: mosaicking and tapering ....... 18
   2.4 Instrumental polarization ...................... 19
   2.5 Ionospheric Faraday rotation ................... 19
   2.6 Missing short spacings ........................ 20

3 The diffuse polarized radio background, and the structure of the Galactic warm ISM 21
   3.1 Introduction .............................. 22
   3.2 Depolarization mechanisms .................... 23
   3.3 The observations ........................... 24
   3.4 The effect of missing short spacings in aperture synthesis observations . 25
   3.5 Depolarization canals ....................... 34
   3.6 Beam depolarization ........................ 40
   3.7 Depth depolarization ........................ 43
   3.8 Relevant components of the ISM ................ 45
   3.9 A model of a Faraday-rotating and synchrotron-emitting layer .......... 47
   3.10 Comparison of model predictions with observations . 53
   3.11 Conclusions ............................. 58
4 Multi-frequency polarimetry of the Galactic radio background around 350 MHz:
   I. A region in Auriga around l = 161°, b = 16° 61
   4.1 Introduction ......................................................... 62
   4.2 The observations ...................................................... 64
   4.3 Total intensity and linear polarization maps .............................. 66
   4.4 Faraday rotation ....................................................... 70
   4.5 The structure in $P$, and the implied properties of the ISM ........... 73
   4.6 Polarized extragalactic point sources .................................. 76
   4.7 Discussion ............................................................ 79
   4.8 Conclusions .......................................................... 83

5 Multi-frequency polarimetry of the Galactic radio background around 350 MHz:
   II. A region in Horologium around l = 137°, b = 7° 87
   5.1 Introduction .......................................................... 88
   5.2 The observations ...................................................... 89
   5.3 Observational results .................................................. 90
   5.4 Observed properties of the ring-like structure .......................... 97
   5.5 The nature of the ring in $P$, $\phi$ and $RM$ ............................ 99
   5.6 Connection of the ring with known ISM objects ....................... 102
   5.7 Structure outside the ring ............................................ 106
   5.8 Conclusions ......................................................... 110

6 Structure in the local Galactic ISM on scales down to 1 pc, from multi-band radio polarization observations 113
   6.1 Introduction .......................................................... 114
   6.2 Distribution of polarized intensity ..................................... 114
   6.3 The nature of the ‘canals’ in polarized intensity ....................... 116
   6.4 The cause of the ‘jumps’ in polarization angle .......................... 117
   6.5 Implications for the structure of the local ISM ....................... 120

7 Polarimetric imaging at 325 MHz of a region in the WENSS survey at 137° <
   $l < 173$° and $-3$° < $b$ < $30$° 121
   7.1 Introduction .......................................................... 121
   7.2 The observations ...................................................... 122
   7.3 Determination of the gradients in polarization angle ................... 127
   7.4 Results ............................................................... 131
   7.5 Large-scale differential rotation measure .............................. 134
   7.6 Conclusions ......................................................... 137

8 Characteristics of the structure in the Galactic polarized radio background at
   350 MHz 139
   8.1 Introduction .......................................................... 139
   8.2 The observations ...................................................... 141
   8.3 Angular power spectrum analysis ....................................... 144
   8.4 Structure functions .................................................... 154
   8.5 Discussion ............................................................ 156
Contents

8.6 Conclusions .................................................. 157

9 MHD simulations of the warm ISM, and comparison with observations:
   preliminary results ........................................... 161
   9.1 Introduction .................................................. 162
   9.2 The numerical simulations .................................. 163
   9.3 Results from the simulations .............................. 166
   9.4 The observations ............................................ 171
   9.5 Comparison of models and observations .................. 174
   9.6 The influence of beam depolarization .................... 176
   9.7 Conclusions ................................................... 180
   9.8 Future work .................................................. 181

Nederlandse samenvatting ........................................ 183

Curriculum vitae .................................................. 197

Nawoord .............................................................. 199
1

Introduction

The enormous and, by human standards, almost entirely empty volume of space in between the stars is in fact filled with a highly intricate interstellar medium (ISM) that consists of many components in complex interaction. The ISM provides the material from which stars form, governs many stellar properties, and, when stars die, is replenished with their processed material.

The interstellar medium of our Galaxy contains gas, dust, comets, meteorites, radiation, cosmic rays, and magnetic fields. The three main components, viz. gas, magnetic fields, and cosmic rays, are approximately in equipartition, i.e. their energy densities are about equal. Therefore, the gaseous ISM, magnetic fields, and cosmic rays cannot be considered in isolation.

The gaseous component of the interstellar medium is by no means a quiet and undisturbed gas. Supernova explosions and strong stellar winds sweep up the gas in filaments, bubbles and shells and cause deformation of magnetic field lines. Shock waves, cosmic rays and Galactic shear input energy into the gas, which is dissipated again in turbulence, line cooling and dust depletion. Coupling of the Galactic magnetic field lines to the interstellar plasma shapes the medium and has a profound influence on star formation.

The structure in the ISM is well characterized by turbulence, i.e. the energy density of the structure can be described by a power law of angular scale, and energy injected at large scales can cascade down to smaller scales. There is an increasing amount of observational support for this description of the ISM, e.g. from turbulence in molecular clouds (Larson 1981), and from the structure of the ionized ISM on parsec scales (Minter and Spangler 1996) down to scales of $10^5$ cm (Rickett 1990, Simonetti et al. 1984).

Theoretical studies also support the description of the structure of the ISM in terms of turbulence, although a turbulent ISM is far too complex to describe analytically. Numerical (magneto-)hydrodynamical simulations have been developed to describe parts of the turbulent nature of the ISM (see reviews by Vázquez-Semadeni 2000, Mac Low 2002, and Zweibel et al. 2002).

The structure of the warm ISM and of the magnetic field in the Milky Way is studied in this thesis by means of polarimetry of the diffuse Galactic radio background. In the introduction, the three major components of the ISM (i.e. gas, magnetic fields and cosmic rays) will first be presented separately, and subsequently the local neighborhood of the Sun. Next, the warm ionized gaseous component of the ISM and the structure of the Galactic magnetic field shall be discussed. Finally, a short outline of the thesis will be given, and a summary of its main results.
<table>
<thead>
<tr>
<th>Component</th>
<th>( T ) (K)</th>
<th>( n ) (cm(^{-3}))</th>
<th>( f ) (%)</th>
</tr>
</thead>
<tbody>
<tr>
<td>Cold molecular</td>
<td>10 - 20</td>
<td>( 10^2 - 10^6 )</td>
<td>( \sim 0.5 )</td>
</tr>
<tr>
<td>Cold atomic</td>
<td>50 - 100</td>
<td>20 - 50</td>
<td>( \sim 0.5 - 5 )</td>
</tr>
<tr>
<td>Warm atomic</td>
<td>6000 - 10000</td>
<td>0.2 - 0.5</td>
<td>( \sim 10 - 50 )</td>
</tr>
<tr>
<td>Warm ionized</td>
<td>( \sim 8000 )</td>
<td>0.2 - 0.5</td>
<td>( \sim 5 - 40 )</td>
</tr>
<tr>
<td>Hot ionized</td>
<td>( \sim 10^6 )</td>
<td>( \sim 0.003 )</td>
<td>( \sim 20 )</td>
</tr>
</tbody>
</table>

**Table 1.1**: Temperature \( T \), density \( n \) and filling factor \( f \) of the gaseous components of the Galactic ISM. Values from Ferrière (2001), Heiles and Kulkarni (1987), and references therein.

### 1.1 The gaseous component of the ISM

About 99% of the mass of the material in the ISM is gas, the remaining 1% is dust. The Galactic gas is present in the form of neutral and ionized, and atomic and molecular species, in three temperature domains of approximately 10 - 100 K, \( 10^4 \) K and \( 10^6 \) K. See Table 1.1 for a brief summary of the characteristics of the gaseous components. The densities of the gas components are such that these three temperature regimes are roughly in pressure equilibrium, but thermally unstable gas at intermediate temperatures has been observed as well (Heiles 2001).

#### 1.1.1 Cold molecular gas

The major part of the gas in the ISM is contained in giant cloud complexes of cold and dense molecular gas. These clouds have temperatures of a few tens of Kelvin and densities of \( 10^2 \) to \( 10^6 \) particles cm\(^{-3}\). With masses of \( \sim 10^5 - 10^7 \) M\(_\odot\), the molecular clouds are the largest coherent objects within the Galaxy. They are confined to the spiral arms. The vertical distribution of molecular clouds has a width of about 80 pc; within this disk, the clouds move with a characteristic turbulent velocity \( v_{\text{rms}} \approx 10 \) km s\(^{-1}\). However, molecular gas has also been detected in high velocity clouds at large distances from the Galactic plane (Aksen and Blitz 1999).

Molecular clouds form by gravitational collapse of diffuse warmer clouds, when more and more effective shielding from heating UV radiation by dust causes the clouds to cool. The clouds fragment, and show structure on many scales, from a few tens of parsec in size to the small dense cores where star formation takes place. Most of the star formation in the Milky Way occurs inside such giant molecular clouds.

#### 1.1.2 Cold and warm atomic gas

Atomic HI gas exists in two temperature ranges which are in thermal equilibrium: the cold neutral medium (CNM) of \( T \approx 100 \) K and \( n \approx 20 - 50 \) cm\(^{-3}\), and warmer diffuse neutral clouds (WNM) with \( T \approx 8000 - 100000 \) K, and \( n \approx 0.01 - 1 \) cm\(^{-3}\). The WNM surrounds the cold CNM clouds, but also pervades much of the Galaxy as an “intercloud medium”. The atomic gas is most readily traced through the famous HI 21-cm spin-flip line, predicted to be detectable in 1944 by Van de Hulst. In studies of HI emission lines, both narrow and broad peaks of the cold and warm components, respectively, can be
detected, while HI absorption studies only yield the cold component (as the absorption coefficient corrected for stimulated emission is inversely proportional to temperature).

Neutral hydrogen is contained in a layer which has an approximately constant thickness of about 200 pc for Galactic radii $0.4R_\odot < R < R_\odot$. At smaller radii, the thickness of the layer decreases to $\sim 100$ pc, and at Galactic radii larger than the Sun’s radius it increases (Lockman 1984, Diplas and Savage 1991). This increase is easily explained by the decreasing gravitational potential in the disk at larger Galactic radii. Although the distribution is flat in the inner Galaxy, there is a systematic warp in the outer Galaxy (Kerr et al. 1957). The neutral gas follows the stellar spiral arms (Oort et al. 1958). There is much small-scale structure in HI, mostly due to supernovae and stellar winds, in the form of e.g. shells, filaments and bubbles. HI clouds have been found many hundreds of parsecs above the Galactic plane. For an extensive review, see Dickey and Lockman (1990).

1.1.3 Warm ionized gas

Two different distributions of the warm ionized medium (WIM) exist: a WIM component around hot stars, and a layer of WIM spread throughout the Galaxy. The concentrations of warm gas around stars are produced through photo-ionization by the intense UV radiation of hot and massive O and B stars in the Galactic disk. A spherical region around the star, the so-called Strömgren sphere or HII region, is ionized out to the distance where the recombination rate equals the ionization rate, typically a fraction of a parsec to a few parsecs. However, most of the warm ionized gas is contained in the second component: an extended, low-density layer centered on the Galactic plane. This so-called Reynolds layer has a height of about a kpc and a filling factor $f \approx 5 - 40\%$ (Reynolds 1991, Kulkarni and Heiles 1987), a pressure equilibrium temperature of about 8000 K and a density comparable to that of the WNM. The mass of the HII is approximately a third of that of the HI, and the power required to maintain the layer is approximately equal to the power input from supernovae. The heating source of the gas is still unknown, but in addition to photo-ionization, another source is needed (Reynolds et al. 1999). Suggested heating sources are photo-electric ejection from dust grains (Reynolds and Cox 1992), dissipation of plasma turbulence (Minter and Spangler 1996), and/or magnetic reconnection (Birk et al. 1998).

1.1.4 Hot ionized gas

The presence of a hot ($T = 10^5$ to $10^7$ K) component (HIM) in the ISM became apparent from interstellar absorption line measurements of highly ionized species (e.g. OVI, NV) and from the detection of the X-ray background. The gas is collisionally ionized by shock waves from supernovae traveling through the medium, and to a lesser extent by strong stellar winds (McGray and Snow 1979, Spitzer 1990). Estimates of the scale height of the HIM from various absorption lines range from $\sim 1.6$ kpc to 5.1 kpc (see review by Ferrière 2001). The filling factor of the HIM is estimated to be $f \approx 17 - 20\%$ (Heiles 1990, Ferrière 1998).
1.2 The local ISM

There is no well-established definition of the "local ISM", but usually the local ISM refers to an irregularly shaped volume of hot gas surrounding the Sun, out to a distance of about 30 pc in the Galactic plane and to 200 pc out of the plane. Observations of the soft X-ray background (Cox and Reynolds 1987) suggest that this volume is a bubble of hot gas with a specific boundary at a varying distance from the Sun, the Local Bubble (LB). The LB is very asymmetric, extending several hundreds of parsecs in the direction perpendicular to the Galactic plane, but very much less in the plane (Snowden et al. 1990). It consists of ionized gas having a temperature of $\sim 10^6$ K and a density of $\sim 0.001$ cm$^{-3}$. The origin of the LB is still uncertain. It could have been created from a single supernova if its ambient density is less than average, or from a multiple supernovae event, or from two supernovae at different times: one supernova to blow the bubble, and another to heat the gas to $10^6$ K. Furthermore, it is not known if the boundary of the LB is still expanding in a shock wave and ionizing new material (which would point to a recent origin $10^8$ or $10^9$ years ago), or if pressure equilibrium has set in and the LB boundary is in a quiescent phase, thereby allowing a much higher age.

Many of the stars in the LB, with distances less than a few tens of parsecs, have hydrogen absorption components in their optical spectra, indicating that cloudlets of diffuse warm hydrogen exist inside the LB (Crutcher 1982, Lallement et al. 1986). The cloudlets have a lower density than the standard diffuse clouds, are a few parsecs in size and are possibly very abundant in the LB (estimated 2000 clouds if they are spherical). The Sun is embedded in one of them, the Local Interstellar Cloud (LIC, Lallement and Bertin 1992), which is also named the Local Fluff due to its low density (Frisch 1998). The LIC is flattened in the Galactic plane, extends mainly in the second quadrant and is about 5 pc in size (Redfield and Linsky 2000). The Sun is positioned at the edge of the LIC ($\sim 30000$ AU), which extends approximately away from the Galactic center and towards the south Galactic pole. The density of the LIC is estimated to be $\sim 0.1$ cm$^{-3}$ and its temperature $T \approx 7000$ K. There are several models for the origin of the diffuse clouds in the LB. The LIC and nearby diffuse clouds could be part of a wall of material in between the LB and the nearby bubble "Loop I" (Brueheiler 1984). Another explanation for some of the very diffuse clouds is that they are expanded remnants of planetary nebulae (Cox and Snowden 1986). The thermal pressure in the LIC ($P/k \approx 1.2 \times 10^{2}$ cm$^{-3}$ K) is about 10 times lower than the thermal pressure of the LB ($P/k \approx 1 - 2 \times 10^{4}$ cm$^{-3}$ K, Bower et al. 1995). This is an unstable situation, and equilibrium requires other sources of pressure, e.g. magnetic fields, if the LIC is a stable pressure-confined cloud.

There are almost no direct observational tracers of the magnetic field inside the LB. However, just outside the LB, Zeeman splitting in cold gas, and starlight polarization due to dust clouds, give possibilities for estimating the influence of the LB on the Galactic magnetic field, see Section 1.3. Heiles (1998) argues that the magnetic field is deformed by the LB, as would be expected if the LB is a bubble blown into the surrounding ISM. For reviews of the local ISM, see e.g. Frisch (1998) or Ferlet (1999).
1.3 The Galactic magnetic field

Beyond the solar system, the magnetic field can only be probed indirectly, and different methods exist to study magnetic fields in different phases of the ISM. These methods are based on polarimetry at optical, infrared or radio wavelengths. For reviews on techniques of measuring the Galactic magnetic field, see Heiles (1976) or Wielebinski and Krause (1990).

Magnetic fields in dust clouds can be studied by measuring the polarization of starlight. In the presence of a magnetic field, elongated dust grains spin perpendicular to the magnetic field, and polarize starlight propagating through the dust cloud (Davis and Greenstein 1951). Optical polarization observations then yield the magnetic field component perpendicular to the line of sight in dust clouds. Several decades ago, it became apparent from extensive surveys that the Galactic magnetic field is aligned along the spiral arms (e.g. Hiltner 1949, Mathewson and Ford 1970).

Magnetic fields cause Zeeman splitting of molecular and atomic spectral lines. Zeeman splitting can be observed at optical and radio wavelengths, in absorption as well as in emission, but because the frequency splitting is so slight, Doppler or thermal broadening can make the Zeeman splitting difficult or impossible to observe. Therefore, the method is used mostly to study magnetic fields in cold molecular clouds (Verschuur 1969, Troland and Heiles 1986).

In the ionized ISM, magnetic fields can be detected through observations of synchrotron radiation. Linearly polarized synchrotron radiation is emitted by cosmic ray electrons spiraling around the magnetic field lines. Therefore, polarimetry yields directly the magnetic field component perpendicular to the line of sight, provided that the observing frequency is sufficiently high that Faraday rotation is negligible. Faraday rotation modulates the synchrotron radiation by rotating the plane of linear polarization, if the radiation propagates through an ionized medium containing a magnetic field. This is due to the birefringence of the magneto-ionized medium, i.e. left and right handed circularly polarized radiation have slightly different phase velocities in a magneto-ionized medium.

The rotation of the polarization angle is proportional to the square of the wavelength of the radiation, and the proportionality constant is defined as the rotation measure \(RM\). The \(RM\) depends on the magnetic field component parallel to the line of sight \(B||\), on the thermal electron density \(n_e\), and on the path length \(ds\) through the medium. The polarization angle \(\phi\) follows from

\[
\phi = \phi_0 + RM \lambda^2 \quad \text{where} \quad RM = 0.81 \int B_\parallel [\mu G] \, n_e [\text{cm}^{-3}] \, ds [\text{pc}]
\]

where \(\phi_0\) is the intrinsic polarization angle at \(\lambda = 0\) (Gardner and Whiteoak 1966, Burn 1966). At high frequencies Faraday rotation is negligible, but at lower frequencies, it is an important effect, which can be used to estimate the strength of the magnetic field along the line of sight. Because the magnitude of Faraday rotation scales with the free electron density, the medium that is primarily probed by this method is the warm ionized ISM (see Table 1.1).

Several sources of polarized radio emission can be used to investigate the Faraday rotation of the polarized radiation by the ISM. Polarized extragalactic point sources
are a valuable tool for the study of the large-scale Galactic magnetic field in the disk and in the halo, due to their uniform distribution across the sky. A second class of sources emitting polarized radiation are pulsars, which have the advantage that, unlike extragalactic sources, they do not possess an intrinsic rotation measure and that their distances are sometimes known. Furthermore, the frequency dependence of group velocity of the pulses traveling through the ISM causes a time dispersion in the pulse signal over frequency, called the dispersion measure $DM = \int n_e \, ds$. So combined $DM$ and $RM$ measurements in the direction of pulsars give a direct estimate of the line of sight component of the Galactic magnetic field, although the measured magnetic field is the integral over all contributions along the line of sight. Finally, observation of linearly polarized diffuse synchrotron emission, as mentioned above, is a powerful probe of Faraday rotation, because it can provide data on rotation measure along many adjacent lines of sight, allowing direct imaging of the electron-density-weighted magnetic field.

In all of these studies, the Galactic magnetic field is treated as composed of a regular component, which follows the spiral arms, and a random component, which varies on small scales. Studies of pulsars and polarized extragalactic sources mainly focus on the regular magnetic field component, which yields information on e.g. the strength and direction of the regular magnetic field, magnetic arms and magnetic dynamos. Diffuse synchrotron emission is an ideal tool for studying the small-scale random magnetic field, allowing e.g. the strength of the random magnetic field component and the role of the magnetic field in turbulence to be estimated. This thesis explores polarimetry of the diffuse synchrotron background in the Galaxy to study the small-scale structure in the ISM and the Galactic magnetic field, and to estimate the global strength and direction of the regular magnetic field component.

Previous Faraday rotation studies using extragalactic sources, pulsars and diffuse synchrotron emission have shown that the Galactic magnetic field is directed approximately along the spiral arms. Combining all available data, the strength of the local Galactic magnetic field is estimated to be around $6 \pm 2 \mu G$ (Han 2001), increasing towards the Galactic center to $\sim 10 \mu G$ at $R = 4$ kpc (Rand and Lyne 1994, Beck et al. 1996).

Vallée (1995) summarized earlier determinations of the pitch angle of the spiral arms, from the direction of the regular Galactic magnetic field from $RMs$ of pulsars and extragalactic point sources and from the spatial location of dust, molecular clouds, $HI$, and $HII$. He concluded that the best estimate for the pitch angle of the spiral arms is $p = -12^\circ \pm 1^\circ$.

In the inner Galaxy, two large-scale reversals in the Galactic magnetic field occur, one between the Local arm and the inward Sagittarius arm, at a Galactic radius of $R \approx 8$ kpc, and the second one further inwards at $R \approx 5.5$ kpc (Rand and Lyne 1994). There are indications both for and against more reversals inside the solar circle, and also for reversals outside the solar circle. For a summary of the controversy see Beck et al. (1996) or Beck (2001).

Estimates of the ratio of the strengths of the regular and the random magnetic field from diffuse polarization are $\sim 1$, with a total field strength $B_t \approx 6 \mu G$ (Berkhuijsen 1971, Phillips et al. 1981). However, studies of pulsar $RMs$ yield $B_{reg} \approx 1.5 \mu G$ and $B_{ran} \approx 5 \mu G$, so $B_{ran}/B_{reg} \approx 3$ (e.g. Rand and Lyne 1994, Han and Qiao 1994). The
$B_{\text{reg}}$ determination from pulsar $RMs$ could be underestimated due to the possibility of still undetected magnetic field reversals. Furthermore, the variations in $B$ and $n_e$ are assumed to be uncorrelated along the line of sight. If $B$ and $n_e$ are anticorrelated, the $B_{\text{reg}}$ estimate is a lower limit, if they are correlated, the $B_{\text{reg}}$ estimate is an upper limit (Beck 2001).

The ratio of regular to random field strengths also seems to vary with position. For example, some spiral galaxies show “magnetic arms” of very regular magnetic field in between the optical arms of the galaxy, while in the optical arms a large $B_{\text{ran}}$ is present, which causes depolarization in the optical spiral arms (Beck and Hoernes 1996). There are indications from $RMs$ of pulsars that this is the case in our own Galaxy as well (Han and Qiao 1994, Indrani and Deshpande 1998).

The random magnetic field has structure on many scales. Single-cell-size models constructed to explain the observed distribution of pulsar $RMs$ yield cell sizes that are typically $\sim 10-100$ pc (Rand and Ulrich 1989, Ohno and Shibata 1993). On the other hand, Minter and Spangler (1996) find that the magnetic field structure can be described by a power law on scales from 0.01 pc to 100 pc. In external galaxies, the characteristic scale of the random magnetic field in the Galactic halo/thick disk seems to be much larger than that in the thin stellar disk, viz. about 100 to 1000 pc (Dumke et al. 1995). For extensive reviews on magnetic fields in all components of the Milky Way, see Vallée (1997, 1998).

1.4 Cosmic rays and synchrotron radiation

Cosmic rays are another important constituent of the ISM, in approximate equipartition with gas and magnetic fields. Cosmic rays comprise mostly protons, $\sim 10\%$ helium nuclei (this fraction increases with cosmic ray energy), $\sim 1\%$ heavier nuclei, and about $1\%$ electrons at several GeV. The ratio of electrons and protons is determined from actual detection in the solar system. It is frequently thought to apply throughout the Galaxy and is often extended to external galaxies and double radio sources.

Typical velocities are close to the speed of light, and the energy ranges from rest energy ($\sim 1$ GeV for protons, $\sim 0.5$ MeV for electrons) to an upper limit of $\sim 3 \times 10^{20}$ eV, where the upper limit may only be induced by small-number statistics (Erlykin and Wipfendale, 2001). The origin of cosmic rays is not completely clear. Low-energy cosmic rays are thought to originate in late-type stars, and to be injected into the medium by coronal flares, or in supernova explosions. They can be accelerated by shocks. There is no consensus about the origin of high-energy cosmic rays, but at least $90\%$ must be of Galactic origin, in view of the dependence of cosmic ray energy on Galactic radius and scale height.

Cosmic ray propagation through the ISM is constrained by the Galactic magnetic field. Cosmic ray particles travel around magnetic field lines, where the spiraling motion of relativistic electrons generates non-thermal synchrotron radiation. The synchrotron emissivity of the Galaxy forms a diffuse background, that was modeled by Beuermann et al. (1985), using the all-sky non-thermal emission maps of Haslam et al. (1981, 1982) at 408 MHz. Their model incorporates the spiral arms, and consists of a thin disk superposed on a thick disk of synchrotron radiation. The half equivalent width
of the thin disk is about 180 pc at the solar radius, that of the thick disk 1800 pc
(with the assumption that the galactocentric radius of the Sun $R_0 = 10$ kpc). Both
equivalent widths increase with galactocentric radius.

The synchrotron emission is Faraday-rotated as well as depolarized while propagating
through the ISM. Multi-frequency polarimetry of the Galactic synchrotron background
therefore enables studies of the physical properties of the ISM.

1.5 ISM in external galaxies

Although studies of the ISM and magnetic fields in external (spiral) galaxies have
inevitably been carried out with a much lower spatial resolution than studies of the
Milky Way, they have the advantage that they delineate the global structure of the
magnetic field. Observations of the synchrotron total intensity give the strength of
the total magnetic field, i.e. both small-scale (turbulent) and large-scale (regular)
components, while the intensity of the linearly polarized emission traces only the regular
magnetic fields (Beck et al. 1999).

Most spiral galaxies show a regular magnetic field component that is aligned with
the optical spiral arms (e.g. Krause (1990) for a review on 7 large nearby face-on spirals, or
Dumke et al. (1995) for 5 edge-on spirals), although magnetic fields perpendicular to
the plane of the galaxy have been observed as well (NGC 4631, Dumke et al. 1995). The
average degree of polarization in the studied external galaxies at a short wavelength of
3 cm (where Faraday rotation is negligible) is 10 to 20%. This degree of polarization
can be explained by a turbulent (or unresolved and tangled regular) magnetic field
component of $B_{\text{ran}}/B_{\text{reg}} = 2 - 3$. However, in the inter-arm regions, the degree of
polarization can be as high as 50%, so that $B_{\text{ran}}/B_{\text{reg}} \approx 0.8$. In the spiral arms, the
average degree of polarization is no more than 0.5 - 5% at $\lambda = 3$ cm, which can be
explained by beam depolarization due to a turbulent field on scales of order 20 pc
(Beck et al. 1999). Dynamo models are able to generate a higher regular magnetic field
component in the inter-arm regions than in the spiral arms (Rohde et al. 1999).

Some spiral galaxies show an axisymmetric magnetic field configuration (i.e. the
magnetic field is in the same direction in each spiral arm), others indicate a bisymmetric
magnetic field (i.e. the magnetic field reverses direction along each spiral arm), but
in many galaxies the configuration of the large-scale magnetic field is unclear. As
mentioned before, the structure of the large-scale magnetic field in our own Galaxy is
also still a matter of debate.

1.6 Polarization as a probe of structure in the warm ISM

In 1987, the first map of the small-scale structure in the linearly polarized synchrotron
background was presented by Junkes et al. (1987). Six years later, Wieringa et al.
(1993) discovered polarization filaments at 325 MHz that did not have a counterpart in
total intensity, and interpreted the small-scale polarization structure as due to Faraday
rotation in a Faraday screen. Since then, many groups have carried out high-resolution
polarimetry (e.g. Duncan et al. 1997, Gray et al. 1998, Duncan et al. 1999, Uyaniker
et al. 1999, Gaensler et al. 2001, this thesis). All of these studies show intriguing
polarization structure, mostly unaccompanied by total intensity structure on the same angular scales. All observations except the ones discussed in this thesis and those done by Wieringa et al. were performed at frequencies of 1.4 GHz and higher to be able to probe the large RMs present in and close to the Galactic plane.

A major survey that provides linear polarization data at 1.4 GHz in a large portion of the Galactic plane is being performed as part of the International Galactic Plane Survey (IGPS). The IGPS consists of the Canadian Galactic Plane Survey (CGPS, Taylor et al. 2002) for a large part of the Northern hemisphere, and the Southern Galactic Plane Survey (SGPS, Dickey et al. 2000) in the Southern hemisphere. The remaining part of the Galactic plane is covered by the VLA (VGP, VLA Galactic Plane Survey), for which only HI line and continuum data are available, but (as yet) no polarization.

The observations presented in this thesis are complementary to the work summarized above for several reasons. First, our observations have been made at 85 cm (350 MHz). Measurements at these long wavelengths are sensitive to much lower rotation measures than the ones probed by studies at frequencies of 1.4 GHz and higher. This implies that 350 MHz observations are more sensitive to depolarization due to propagation through large volumes of magneto-ionic medium as well, causing an effective "polarization horizon" of about 500 - 1000 pc. So we only probe the closest regions of the warm ISM. Secondly, the fields that we study are at intermediate to high latitudes. At long wavelengths, we are able to study these regions of low density, and thereby gain insight into the properties of the warm ionized ISM at higher Galactic latitudes.

1.7 Outline of the thesis

The research described in this thesis uses the Westerbork Synthesis Radio Telescope (WSRT) in the Netherlands. Two regions were selected on the basis of polarization maps produced as a by-product of the Westerbork Northern Sky Survey (WENSS, Rengelman et al. 1997), a radio survey of all declinations above 30° at 325 MHz. These two regions showed such intriguing polarization structure that it was decided to reobserve them with greater sensitivity and in multi-frequency mode. The regions are each approximately 60 square degrees in size and located at intermediate and high latitudes, in the second Galactic quadrant. In these directions, the regular Galactic magnetic field is almost perpendicular to the line of sight. Both fields display an abundance of structure in polarized intensity and in polarization angle, that has no correlated structure in total intensity. We analyze these observations, and compare them to numerical simulations. In addition, we present polarization maps of a part of the WENSS survey itself.

In Chapter 2, we briefly discuss the (polarization) characteristics of the WSRT, and its strengths and weaknesses for this work.

In Chapter 3, we show that polarization data suffer from several depolarization effects. To correctly interpret the data, a careful analysis of the imprints of depolarization and of the WSRT are necessary. A study of different depolarization processes (beam depolarization and depth depolarization) and the influence of missing short spacings of the interferometer is therefore performed in this chapter. We conclude from the
width of the distribution of rotation measure \( RM \) and from the good quality of the linear relation between polarization angle \( \phi \) and \( \lambda^2 \), that missing large scale structure in Stokes \( Q \) and \( U \) cannot play a major rôle in the observations discussed in this thesis. Secondly, beam depolarization is important in narrow depolarization “canals” of width \( \sim 1 \) beam size. The propagation of radiation through a model medium containing thermal electrons, cosmic rays and magnetic field is considered. Radiation emitted in and propagating through this medium is Faraday-rotated and depolarized. By comparing “observables” from the model, such as rotation measure and polarized intensity, with observables from the WSRT observations, conclusions about properties of the ISM are drawn. The model indicates that the regular Galactic magnetic field component is at least as strong as the random magnetic field component. Strong regular magnetic field is expected in the inter-arm region where the probed magnetic field is located. Furthermore, our two fields were selected on the basis of their unusually strong structure in \( P \), which selects a high regular magnetic field strength as well. Structure in the random component of the magnetic field on scales of 10 - 20 pc can cause the observed degree of depolarization and rotation measure distribution. Finally, the distance of the “polarization horizon”, i.e. the distance beyond which most polarized emission is depolarized, is estimated to be approximately 600 pc.

A detailed discussion of the observations of the two regions is presented in Chapters 4 and 5. The first region, discussed in Chapter 4, shows filaments of high polarized intensity on typical scales of degrees, where the Galactic magnetic field and thermal electron density must be homogeneously distributed. Furthermore, the observations are consistent with structure in the random component of the magnetic field on a characteristic scale of about 20 pc, while the regular magnetic field is slightly larger than the random component.

In Chapter 5, the observations of the second region are discussed, in which there is a ring-like structure in polarized intensity with a radius of about 1.4'. In polarization angle and rotation measure, there is no ring but rather a filled, disk-like structure with a radius of \( \sim 1.7' \). This structure must be immersed in an extremely homogeneous environment, over the entire line of sight. Several possible explanations for the ring-like structure are discussed. We conclude that the structure is due to a local reversal in the magnetic field direction, possibly accompanied by enhanced thermal electron density. We propose a funnel-like structure for the magnetic field directed away from the observer.

In Chapter 6, we describe a property of the polarized intensity which is observed in all fields, namely narrow one-dimensional structures where the polarized intensity decreases and becomes negligible, so-called depolarization canals. From data in the region discussed in Chapter 4, we conclude that depolarization canals are due to abrupt changes of 90° \( \pm n \times 180° \) in polarization angle. The changes in polarization angle are most likely caused by abrupt changes in \( RM \) of the right magnitude to produce \( \Delta \phi = 90° \pm n \times 180° \). Therefore, the depolarization canals are not narrow ridges in the medium, but rather a boundary between two regions of different \( RM \).

We present diffuse polarization maps from the Westerbork Northern Sky Survey (WENSS, Rengelink et al. 1997) in Chapter 7. The WENSS survey is a radio survey performed with the WSRT at 325 MHz with about 1' resolution. Although the WENSS covered the major part of the northern hemisphere, only a subset of the WENSS fields...
can be used for polarization analysis, viz. those that were observed entirely or mostly at night. This is because the ionospheric disturbance during sunrise and sunset, and the interference from the Sun itself during the day, preclude the study of the faint diffuse polarized emission from the Galaxy. Since the WENSS is a single-frequency survey, it cannot be used to obtain rotation measure information, but it provides a large-scale map of Galactic polarization, from the Galactic plane to high Galactic latitudes, and covers a large part of the second Galactic quadrant. Structure in polarization angle is clearly aligned with Galactic longitude, as could be produced by a gradient in $RM$ perpendicular to the Galactic plane.

In Chapter 8, we give a statistical analysis of the turbulent nature of the ISM in the two regions described in Chapters 4 and 5, and in the polarization maps from the WENSS survey, discussed in Chapter 7. Power spectra and structure functions of rotation measure, linearly polarized intensity and Stokes parameters $Q$ and $U$ are computed, discussed and compared with previous studies. The power spectra can be well described by a power law over the available spatial scales of $\sim 0.1^\circ$ to $\sim 1^\circ$. The power spectrum of $RM$ has a logarithmic slope $\sim -1$, and the slope of the spectrum of $P$ is $\sim -2.2$. The slopes of power spectra of $Q$ and $U$ are even steeper, which suggests the presence of a Faraday foreground screen.

In Chapter 9, preliminary results of three-dimensional magneto-hydrodynamical numerical simulations are described. These simulations model a turbulent medium consisting of warm gas threaded with magnetic fields. The medium acts as a Faraday screen for radiation propagating through it. We analyze polarization at different wavelengths, which propagates through the medium, and use different smoothings of the resulting Stokes $Q$ and $U$ maps to simulate a finite telescope beam. We discuss the influence of our finite observing beam on the determination of rotation measure, and on the power spectra of polarized intensity and rotation measure.

1.8 Main results

From our model of depth depolarization in a magneto-ionic ISM, combined with the observations of our two fields in the second Galactic quadrant at intermediate Galactic latitudes, we conclude that the regular Galactic magnetic field must be directed almost perpendicular to the line of sight in both regions. We also conclude that the regular component of the Galactic magnetic field must be larger than or equal to the random component of the magnetic field. A regular magnetic field component which is larger than the random component is derived in inter-arm regions, in our own Galaxy from pulsar $RM$ observations, and in external galaxies from diffuse synchrotron radiation. This is consistent with our conclusion that the observed polarized emission originates in the inter-arm region between the Local Arm and the Perseus arm, and that emission from the Perseus arm or beyond is completely depolarized.

Intriguing structures in polarization angle and polarized intensity on scales of degrees are observed. In one of the fields, we detect filaments of high polarized intensity many degrees long, which indicate the presence of filaments or sheets of regular Galactic magnetic field, probably in a region of low electron density. The other field shows a circular structure in polarized intensity and polarization angle, which we argue is a
funnel-shaped magnetic field structure, observed along its long axis.

Elongated, beam-sized wide canals of low polarization are present in all observations. We conclude that these canals are likely to be caused by beam depolarization, and are boundaries between regions of constant polarization angle, where the difference in angle across a canal is $90^\circ$. The $90^\circ$ change in angle is most likely caused by an abrupt change in $R M$, but can also be caused by depolarization along the line of sight in a uniform magneto-ionic medium.

We present single-frequency polarization maps over 1000 square degrees of sky in the second quadrant, which show global alignment of polarization angle parallel to the Galactic plane. A gradient in angle over Galactic longitude and Galactic latitude can be derived, which can be interpreted as a gradient in $R M$. Power spectra analysis of subregions in this large field shows that the multipole spectral index decreases with Galactic latitude, i.e. the power spectra of $R M$ become flatter towards higher latitudes.

Comparison of 3-dimensional magnetohydrodynamic simulations of the warm ionized component of the ISM to the observations provides a promising tool to study the propagation of polarized radiation through the ISM. Beam depolarization has a profound influence on the determination of $R M$. However, higher resolution is needed for quantitative conclusions.

References

Beck R., 2001, SSRv 99, 243
Beck R., & Hoernes P., 1996, Nat 379, 47
Bruhweiler F. C., 1984, in Local Interstellar Medium, International Astronomical Union Colloquium No. 81, ed. by Y. Kondo, F. C. Bruhweiler, B. D. Savage, p. 39
Cox D. P., & Snowden S. L., 1986, AdSpR 6, 97
Introduction

Ferrière K. M., 2001, RvMP 73, 1031
Gardner F. F., & Whiteoak J. B., 1966, ARAA 4, 245
Heiles C., 1998, LNP 506, 229
Heiles C., 1976, ARA&A 14, 1
Indrani C., & Deshpande A. A., 1998, NewA 4, 331
Kerr F. J., Hindman J. V., & Carpenter M. S., 1957, Nat 180, 677
Krause M., 1990, in Galactic and intergalactic magnetic fields; Proceedings of the 140th Symposium of IAU, p. 396
Mac Low, M.-M., 2002, in Proceedings of the workshop on Simulations of MHD Turbulence in Astrophysics, ed. by E. Falgarone and T. Passot
McCann D., & Sanders W. T., 1990, ARA&A 28, 657
McCray R., & Snow T. P. Jr., 1979, ARA&A 17, 213
Spitzer L. Jr., 1990, ARA&A 28, 71
Taylor A. R., Gibson S. J., Peracaula M., Martin P. G., Landecker T. L., Brunt C. M.,
Dewdney P. E., Dougherty S. M., Gray A. D., Higgs L. A., Kerton C. R., Knee L. B. G.,
AJ, submitted
Vallee J. P., 1998, FCPH 19, 319
Vallee J. P., 1997, FCPH 19, 1
work Workshop, ed. by V. G. Gurzadyan and R. Ruffini, p.379
Wielebinski R., & Krause F., 1990, A&AR 4, 449
Zweibel E. G., Heitsch F., & Fan Y., 2002, in Proceedings of the workshop on Simulations of
MHD Turbulence in Astrophysics, ed. by E. Falgarone and T. Passot
2

WSRT Polarimetry

2.1 Observing with the WSRT

The Westerbork Synthesis Radio Telescope (WSRT) is an east-west array, consisting of fourteen 25m dishes, four of which are movable. The maximum baseline of the array is 2.7 km. In one 12 hour synthesis, a regular \((u,v)\)-coverage is obtained with a minimal baseline increment of 72m, which gives a first grating ring at a distance from the pointing center of 44' in right ascension, and 44'/\sin \delta in declination at 325 MHz. In this analysis, we combined six 12hr synthesis observations. This yields a minimal baseline increment of 12m, which produces a first grating ring at 4.4' from the pointing center at these frequencies.

The long wavelength front end of each dish contains two orthogonal dipoles \(X\) and \(Y\) mounted in the primary focus, with fixed position angles \(\phi_X = 90^\circ\) (east) and \(\phi_Y = 180^\circ\) (south). Four cross-correlations of the \(X\) and \(Y\) dipoles in each telescope, viz. \(XX\), \(YY\), \(XY\) and \(YX\), allow determination of the full set of Stokes parameters \(I\), \(Q\), \(U\), and \(V\). For the \(XX\) and \(YY\) correlations, all 91 baselines were included, whereas for the correlations \(XY\) and \(YX\), only data from the 40 standard (fixed-movable) baselines were used.

The observations were carried out using the DCB back end, which is a broadband back end with a maximum total bandwidth of 80 MHz. The DCB provides 8 frequency bands of 5 or 10 MHz, the central frequency of which can be set independently. In the observations discussed here, we used 8 frequency bands of 5 MHz wide, centered between 310 MHz and 390 MHz. Due to radio interference and hardware problems, we could use only 5 bands, viz. those centered at 341, 349, 355, 360 and 375 MHz (~90 cm).

Standard calibration techniques were used to calibrate the data, using the newstar (Netherlands East-West Synthesis Telescope Array Reduction) data reduction package, written by W. Brow. newstar was designed for the WSRT, and can only be used for east-west array data. However, newstar optimally uses the many WSRT redundant (i.e. non-standard) baselines, and can create sophisticated source models, using extended sources and containing polarization and spectral index information. We refer to Wieringa (1991) for an extended discussion of calibration of synthesis arrays.

The data were obtained during night time for two reasons. First, a highly polarized response from the sun is picked up through distant sidelobes, which can thoroughly corrupt all data on baselines smaller than about 100m, i.e. on scales above about 0.5'. As we expect diffuse polarization on these scales, it is extremely important to have
good data on the short baselines. Secondly, the Faraday rotation characteristics of the ionosphere change strongly during twilight, which severely complicates the analysis of polarized radiation.

2.2 Calibration of polarization data

The complex visibility which is measured with an interferometer consisting of two dipoles with position angles $\phi_1$ and $\phi_2$ is

\[
v_{12} = \begin{bmatrix} v_{12} \\
\bar{I} \\
\bar{Q} \\
\bar{U} \\
\bar{V} \end{bmatrix} = \begin{bmatrix} G_{12} \\
\tilde{I} \\
\tilde{Q} \\
\tilde{U} \\
\tilde{V} \end{bmatrix} \begin{bmatrix} \cos(\phi_1 - \phi_2) - \epsilon_{12} \sin(\phi_1 - \phi_2) \\
\cos(\phi_1 + \phi_2) - \eta_{12} \sin(\phi_1 + \phi_2) \\
\sin(\phi_1 + \phi_2) + \eta_{12} \cos(\phi_1 + \phi_2) \\
\sin(\phi_1 - \phi_2) + \epsilon_{12} \cos(\phi_1 - \phi_2) \end{bmatrix}
\]

(Weiler 1973), where $\bar{I}, \bar{Q}, \bar{U},$ and $\bar{V}$ are the Fourier transforms of the Stokes parameters $I, Q, U,$ and $V$. Eq. (2.1) describes the visibilities $v$ for all four cross-correlations $XX, YY, XY$ and $YX$ of the two dipoles 1 and 2. The complex gain $G_{12}$ and the $\epsilon_{12}$ and $\eta_{12}$ factors contain the telescope errors. Calibration of the data yields an estimate of these errors, so that they can be corrected for, and the values of the Stokes parameters can be derived from Eq. (2.1). The telescope errors consist of

1. gain error $g$, which is a fractional contribution to the measured intensity. It is caused electronically or in the atmosphere.
2. phase error $p$, which is an additional phase difference between the two correlated dipoles. This can have an electronic and atmospheric origin.
3. dipole error $\Delta$, which is the error in the positioning of the dipoles with respect to the sky. This is a mechanical error, and is usually stable over the length of the observation.
4. ellipticity error $\theta$, which is the deviation of orthogonality between the two correlated dipoles, and has a mechanical origin as well.

These four errors are incorporated in the $G_{12}, \epsilon_{12}$ and $\eta_{12}$ factors. If $\Delta$ and $\theta$ are small, second-order terms can be ignored, and the following relations hold:

\[
\begin{align*}
\epsilon_{12} &= (\Delta_1 - \Delta_2) - i(\theta_1 + \theta_2) \\
\eta_{12} &= (\Delta_1 + \Delta_2) - i(\theta_1 - \theta_2) \\
G_{12} &= g_1g_2 e^{-i(p_1 - p_2)}
\end{align*}
\]

Again ignoring second-order terms, the Stokes parameters $I, Q, U$ and $V$ can be derived from the measured visibilities $v$ in the auto- and cross-correlations $XX, YY, XY$ and $YX$ as

\[
\begin{align*}
\nu_{XX} &= G_{XX} (I + Q) \\
\nu_{YY} &= G_{YY} (I - Q) \\
\nu_{XY} &= G_{XY} (U + iV + \epsilon_{XY} I) \\
\nu_{YX} &= G_{YX} (U - iV - \epsilon_{YX} I)
\end{align*}
\]
(Hamaker et al. 1996, Sault et al. 1996). Using observations of an unpolarized strong calibrator source, Eqs. (2.5) to (2.8) can be solved for complex gains $G_{12}$ and leakage $\epsilon_{12}$. This is the basis of the total intensity calibration, discussed in Section 2.2.1.

Calibration is needed for three additional free parameters for polarized data. The leakage term $\epsilon_{12}$ only depends on the differential dipole angle error and ellipticity, leaving the absolute values as two free parameters. Another degree of freedom is the phase relation between the $X$ and $Y$ gains. These three degrees of freedom can be resolved using a polarized calibrator source, as explained in Section 2.2.2.

The standard calibrator source for the WSRT is 3C286. The absolute flux scale at 325 MHz is based on a value of 26.93 Jy for 3C286, which is the Baars et al. (1977) value. From this value the flux scales of the calibrator sources used for these observations were derived, which are the unpolarized calibrators 3C48, 3C147, and 3C295, and the polarized calibrators 3C345 and 3C303.

### 2.2.1 Total intensity calibration

The most common source of radio interference corrupting the data is man-made interference. In the case of the WSRT this is mostly interference by TV-stations, military communication, and computers in the service building next to the telescopes.

Data of calibrator sources as well as field data were checked for corruption by radio interference and obviously corrupted data were manually flagged. Furthermore, flagging was applied to all data with high Stokes V values. Extra care was taken for short baselines, because those show higher visibilities than the long baselines due to the extended diffuse radiation.

Subsequently, a strong unpolarized calibrator was used to determine the complex gain factors $G_{12}$. By measuring the complex visibilities $v_{XX}$ and $v_{YY}$ for an unpolarized calibrator, one can solve Eqs. (2.5) and (2.6) to determine independent estimates for the gain factors $g_X$ and $g_Y$ and the phase differences $(p_{X,1} - p_{X,2})$ and $(p_{Y,1} - p_{Y,2})$ according to Eq. 2.4. The corrections obtained from the calibrator source are applied to the field data.

Finally, self-calibration is applied. In this procedure, a point source model of the field is made, which is used to self-calibrate the field. The model is built up iteratively, where in each step only the strongest sources are included, to avoid that strong grating rings are counted as sources. After each step, the current model is subtracted and new sources are sought. The model allows for inclusion of extended sources and spectral index information for each source.

### 2.2.2 Polarization calibration

To obtain reliable polarization data, the leakage terms $\epsilon$ in Eqs. (2.7) and (2.8) have to be determined. The leakage consists of dipole angle errors $\Delta$, i.e. errors in the positioning of the dipole, and of ellipticity errors $\theta$, i.e. errors in the orthogonality of the dipoles. They can be determined using a strong unpolarized calibrator source. Incorporating the gain and phase errors determined in the total intensity calibration, one can solve Eqs. (2.7) and (2.8), assuming that $U = V = 0$. This yields a determination of $(\Delta_1 - \Delta_2)$ and $(\theta_1 - \theta_2)$ for both $XY$ and $YX$. 

In the total intensity calibration, only the phase differences \((p_1 - p_2)\) in Eq. (2.4) are calibrated. This means that two arbitrary phase-zero terms \(p_{X,0}\) and \(p_{Y,0}\) are still unknown. The phase-zero terms cancel in the \(G_{XX}\) and \(G_{YY}\) gains, e.g.

\[
G_{XX,12} = g_{x,1}g_{x,2} \exp[-i(p_{x,0} + p_{x,1} - (p_{x,0} + p_{x,2})]]
\]

and similarly for \(G_{YY}\). However, in the determination of \(G_{XY}\) and \(G_{YX}\), a term 

\[\pm(p_{X,0} - p_{Y,0})\]

remains. This is an additional phase-zero difference (PZD), which can corrupt \(U\) and \(V\) data.

The PZD term can be determined using a linearly polarized calibrator of known polarization. Again from Eqs. (2.7) and (2.8), two independent determinations of Stokes parameter \(U\) can be made, assuming \(V = 0\) and taking gain and phase errors into account. The two independent \(U\) values yield the PZD error. However, if the \(U\) signal in the polarized calibrator source happens to be weak (all linear polarization in \(Q\)), this method may not yield the correct phase. This will be visible in the maps of the diffuse data as structure in \(V\) that should have been in \(U\). A practical solution to this, which we had to apply a number of times, is to add manually an additional phase difference until the \(V\)-maps are empty. These corrections are applied to the field data as well.

Self-calibration was applied for the Stokes \(Q\) and \(U\) maps for parts of the data for which no usable calibrator data was available. Models including polarization information were constructed in a similar way as described above.

#### 2.3 Constructing maps: mosaicking and tapering

As the desired field of view is much larger than the primary beam of the telescope, we used the mosaicking technique (e.g. Rengelink et al. 1997). In this observing mode, the array cycles through a preselected set of pointing positions a number of times during the 12 hour observation period. The beam is directed at each pointing position for a certain multiple of 10 seconds before moving to the next pointing center. This is repeated during the 12hr observing period, to produce a constant and sufficient \((u, v)\)-coverage for each of the pointings. Uniform sensitivity is achieved for a distance between pointing positions equal to half of the half-power beam width of the individual telescopes \((\text{HPBW} \approx 2\degr\) for 25m dishes at 350 MHz). This observing mode reduces the instrumental polarization to at most 1%, except in the outer edges of the outermost pointings where only data of a single pointing are available.

In the two multi-frequency observations discussed in this thesis, the number of pointings is 5 × 7 and 5 × 5, respectively. To obtain sufficient sampling of the \((u, v)\)-plane, the time of observation per pointing (dwell time) was 40 and 60 seconds for the two fields, allowing the pointing centers to be observed 31 and 29 times per 12hr observation, respectively. The dwell time includes 10 seconds to move the telescope to the next pointing, and the data in this interval were flagged. The distance between pointing centers was chosen as 1.25\degr, which resulted in two fields of \(9\degr \times 11\degr\) and \(9\degr \times 9\degr\) respectively. As the instrumental polarization increases rapidly at the edges of the fields, we discarded all data approximately a degree from the edges of the fields.
A taper can be used to decrease the resolution of the observations, and so increase the signal-to-noise ratio for extended emission. The taper is applied in the \((u,v)\)-plane by giving long baselines less weight according to a specific tapering function. In these observations, we use a taper with a Gaussian decrease in weight with baseline length. The Gaussian has a value of 0.25 at a baseline of \(\sim 300\) m, where the exact baseline value is chosen so that the resulting resolution of the observations is equal at all frequencies.

From the calibrated, mosaicked and tapered Stokes \(Q\) and \(U\) maps, polarized intensity \(P\) and polarization angle \(\phi\) can be derived as

\[
P = \sqrt{Q^2 + U^2}
\]

\[
\phi = \frac{1}{2} \arctan \left( \frac{U}{Q} \right)
\]

2.4 Instrumental polarization

There are several sources of instrumental polarization, which can be divided in on-axis and off-axis instrumental polarization.

On-axis instrumental polarization is created if the gain calibration of the data is not perfect. In this case, for an unpolarized source, \(v_{XX} \neq v_{YY}\) and therefore \(Q \neq 0\). In the data, this is visible as a non-zero \(Q\) signal of the same sign in all extragalactic sources in the field. This can be corrected for by calculating the magnitude of the instrumental \(Q\) signal. This yields the correction factor for the \(G_{XX}\) and \(G_{YY}\) to make a proper calibration. Also, imperfect calibration of the position and angle of the orthogonal dipoles, or incomplete phase-zero calibration, can cause a small amount of on-axis instrumental polarization. In addition, on-axis instrumental polarization is induced in every telescope because of polarization sensitive reflections. However, this contribution is too small to be relevant for our observations.

Off-axis instrumental polarization induces a quadrupolar instrumental polarization, visible as a “clover-leaf” pattern in \(Q\) and \(U\), in which the “leaves” have alternating sign. This probably originates in the four-legged support structure of the WSRT front ends. The off-axis polarization can reach values of several tens of percent at distances beyond the 20% power point of the primary beam. However, the primary beam weighted combination of the various pointing centers in the rectangular grid gives very low weight to those parts of the primary beam. The clover-leaf polarization pattern thus results in a remaining off-axis polarization that is never more than about 1% in all parts of the mosaic where 4 pointing centers effectively contribute.

2.5 Ionospheric Faraday rotation

The ionosphere, roughly 100 to 1000 km above the earth’s surface, contains ionized particles due to solar radiation and the solar wind. In interaction with the earth’s magnetic field, the free electrons in the ionosphere cause Faraday rotation. The ionospheric Faraday rotation is larger during the day than during the night, because the ionosphere extends further out when the sun heats it directly. Furthermore, ionospheric Faraday rotation is higher in summer than in winter, and higher during solar maximum than during solar minimum.
The contribution of the ionosphere to the measured rotation measures at the location of the WSRT can be estimated from measurements of the total electron content (TEC) of the ionosphere with an ionosonde. In addition, estimates can be made with the “Parametrized Ionospheric Model” (PIM), a global model of theoretical ionospheric behavior.

All observations discussed in this thesis were done during night time, minimizing ionospheric Faraday rotation. For the two multi-frequency observations, no ionospheric corrections were applied, see Chapters 4 and 5. In the WENSS polarization maps, differential ionospheric Faraday rotation was corrected with polarized point sources that are present in the maps, see Chapter 7.

2.6 Missing short spacings

All interferometers have a non-zero minimum baseline, which induces an incomplete coverage of the \((u, v)\)-plane. Therefore, the synthesis telescope is insensitive to large-scale structure. For the WSRT, the maximum angular scale which can be observed without reduced sensitivity is \(\sim 1^\circ\) at 350 MHz. An obvious way to correct this deficit is adding absolutely calibrated single-dish data to the interferometer data, from a single-dish telescope which has baselines overlapping with the interferometer’s (e.g. Stanimirovic 2002). However, this is not possible for the WSRT at 350 MHz, simply because no absolutely calibrated single-dish which is large enough to cover all missing spacings operates at this frequency. As we cannot correct for the missing short spacings, we have to estimate their importance and take them into account in interpreting the data. This is explained in detail in Chapter 3.

References

Wieringa M. H., 1991, PhD thesis Leiden University
The diffuse polarized radio background, and the structure of the Galactic warm ISM

M. Haverkorn, P. Katgert and A. G. de Bruyn

Abstract

Observed structure in the diffuse polarized radio background, uncorrelated to structure in total intensity, must be caused by Faraday rotation, depolarization or instrumental effects. We discuss each of these effects separately and estimate their importance using the results of multi-frequency polarimetric observations of the diffuse background. We use data for two fields of observation located in the second Galactic quadrant at positive latitudes, each more than 50 square degrees, observed with the Westerbork Synthesis Radio Telescope at 5 frequencies near 350 MHz. We show that large-scale components that are missing due to the insensitivity of the interferometer to scales larger than a degree cannot be important in these fields. Beam depolarization (caused by structure in polarization angle within one telescope beam) can only create structure on scales of a beam width. Beam depolarization is important and most conspicuously present in depolarization channels, which denote boundaries between regions of different rotation measure. We conclude that the dominant effect that produces the observed polarized structure is depth depolarization, i.e. depolarization along the path through a medium which emits synchrotron radiation and causes Faraday modulation. A simple single-cell-size model is constructed, and used to predict distributions of observable parameters. Comparison with the observations of the two fields then yields estimates for the allowed ranges of the parameters in the model. Model parameter estimates agree for the two fields of observation. The resulting estimate for the cell size in the thin disk is $15 \pm 10$ pc, with a ratio of random to regular magnetic field $B_{\text{rms}}/B_{\parallel} \approx 0.7 \pm 0.5$. The regular magnetic field component perpendicular to the line of sight is much higher than the parallel component, indicating a regular magnetic field almost perpendicular to the line of sight. We conclude that the distance beyond which most of the polarized emission is depolarized is about 500 to 700 pc, although this is not a well-defined distance.
3.1 Introduction

The omnipresent cosmic rays in the Milky Way, circling in the Galactic magnetic field, provide a synchrotron radio background which is partly polarized. This radiation propagates through the warm ionized interstellar medium (ISM) and is modulated by it. This makes observations of the polarized continuum radio background a valuable tool for studies of the warm ISM.

The first measurements of the diffuse polarized radio background were made with the Dwingeloo 25m dish at 408 MHz by Seeger and Westerhout (1961), and shortly afterwards Bingham and Shakeshaw (1967) performed the first multi-frequency study. They used three frequencies 408 MHz, 610 MHz and 1407 MHz at a resolution of about 2' to study Faraday rotation in the warm ISM. Rotation measures ($RM$) of up to $\sim 8\, \text{rad m}^{-2}$ were obtained in several regions of high polarization. From the measured intrinsic polarization angle, they concluded that the regular magnetic field must be directed along the Galactic plane. Brouw and Spoelstra (1976) discussed several radio surveys at 5 frequencies from 408 MHz to 1411 MHz, which enabled Spoelstra (1984) to determine $RM$ maps of a major part of the northern hemisphere. These maps show structure on many scales, with values of $|RM| \leq 10\, \text{rad m}^{-2}$. However, these $RM$ values are uncertain because the surveys were not fully sampled.

Junkes et al. (1987) published the first maps of sub-degree-scale structure in the polarized diffuse background at $\lambda = 11\, \text{cm}$. However, the first analysis and interpretation of small-scale diffuse polarization was done by Wieringa et al. (1993), who presented observations of discrete structures in polarization that were not observed in total intensity, and interpreted these as Faraday rotation structures. Since the recognition of polarization structures as caused by Faraday rotation, many high-resolution studies of the diffuse polarized background have been done, yielding many unexpected results. Maps of polarized intensity $P$ as well as polarization angle $\phi$ show without exception ubiquitous structure on arcminute to degree scales, mostly uncorrelated with structure in total intensity $I$. Arcminute resolution polarization measurements have revealed formerly unknown HII regions (Gaensler et al. 2001), associated halos (Gray et al. 1999) and depolarizing outflow regions (Duncan et al. 1999). Smooth, symmetrical polarization structures not visible in any other way have been discovered (Gray et al. 1998, Uyaniker and Landecker 2002). In recent years, polarized Galactic foregrounds have also become of interest for cosmology, because they form a possible foreground contamination for polarization measurements of the Cosmic Microwave Background Radiation (CMBR, e.g. Prunet et al. 2000).

In general, there is no significant correlation between the small-scale structure in polarized intensity $P$ or polarization angle $\phi$ and the structure in total intensity $I$. Therefore, the structure in polarization is not believed to be due solely to intrinsic structure in synchrotron emission. Instead, the structure in polarization angle is explained in terms of Faraday rotation of the synchrotron radiation that impinges on the magneto-ionic medium of the ISM relatively close to the Sun (Burn 1966). Multi-frequency polarimetry of the synchrotron emission yields Faraday rotation measures ($RM$), which depend on electron density and magnetic field, and therefore enable the study of the electron-density-weighted strength and structure of the Galactic magnetic field.
However, whereas the hypothesis of Faraday modulation can explain the variation in polarization angle, it does not provide an explanation for the structure in polarized intensity $P$. Although the lack of zero-baseline visibilities in some interferometric observations could produce structure in $P$ from structure in $\phi$, this would leave the structure in $P$ in absolutely calibrated single-dish observations unaccounted for. In this paper, we discuss the various processes that can produce structure in $P$, and we use several multi-frequency datasets obtained with the Westerbork Synthesis Radio Telescope (WSRT) to gauge their importance. We estimate the influence of missing short spacings on the polarization data, and of several depolarization mechanisms. A simple model simulating depolarization processes in a magnetoionic medium is compared to the results of radio polarimetric observations. This allows an estimation of the random and regular components of the Galactic magnetic field, the correlation length of the random magnetic field, and the distance out to which polarization can be observed.

The paper is organized as follows. We discuss the different depolarization mechanisms and nomenclature used in the literature in Section 3.2. In Section 3.3 we summarize the relevant parameters of the Westerbork polarization observations that will be used to estimate the importance of the various effects that contribute to the structure in $P$. In Section 3.4 the role of missing short spacings in interferometer measurements is estimated, and we discuss how the resulting images can be affected. Section 3.5 presents a discussion on the origin of depolarization canals in polarized intensity. In Section 3.6 beam depolarization is discussed, and Section 3.7 describes depth depolarization in a layer that contains both synchrotron-emitting and Faraday-rotating material. In Section 3.8 we discuss the components of the ISM relevant to the model of depth depolarization that is presented in Section 3.9. The predictions of the model calculations are compared to the observations in Section 3.10, which yields a statistical description of the properties of the emitting and Faraday-modulating ISM. Finally, we present a discussion and conclusions in Section 3.11.

### 3.2 Depolarization mechanisms

Two different approaches to the description of depolarization mechanisms can be found in the literature: one distinguishing the mechanisms according to the physical processes that cause the depolarization, the other is a purely geometrical division. In this paper, we use the latter, which is more convenient for our purposes. We will first describe the physical processes causing depolarization, and then how these are treated here. For extensive treatments of the depolarization processes see Gardner and Whiteoak (1966), Burn (1966), Sokoloff et al. (1998) and Fletcher (in prep.).

- **Wavelength independent depolarization** is the depolarization due to turbulent magnetic fields in the ISM. Turbulent magnetic fields emit synchrotron radiation with varying polarization angle. Therefore, superposition of the polarized radiation along the line of sight and across the telescope beam results in partial depolarization of the emission, independent of wavelength. No Faraday rotation is involved.
Table 3.1: WSRT polarization observations in the constellations Auriga and Horologium. Given are position and size of the regions, observed frequency bands and number of pointings used to mosaic the region.

<table>
<thead>
<tr>
<th>Constellation</th>
<th>Position</th>
<th>Size</th>
<th>Frequency (MHz)</th>
<th>Resolution</th>
<th>Pointings</th>
</tr>
</thead>
<tbody>
<tr>
<td>Auriga</td>
<td>(161°, 16°)</td>
<td>7°×9°</td>
<td>341, 349, 355, 360, 375</td>
<td>5.0'×6.3'</td>
<td>5×7</td>
</tr>
<tr>
<td>Horologium</td>
<td>(137°, 7°)</td>
<td>7°×7°</td>
<td>341, 349, 355, 360, 375</td>
<td>5.0'×5.3'</td>
<td>5×5</td>
</tr>
</tbody>
</table>

- **Differential Faraday rotation** occurs if a uniform medium contains thermal and relativistic electrons and magnetic field. Synchrotron radiation emitted at different distances along the line of sight will undergo a different amount of Faraday rotation along its path length. This is a one-dimensional, wavelength dependent depolarization effect.

- **Internal Faraday dispersion** is the depolarization in a turbulent synchrotron-emitting magneto-ionic medium. Variation in intrinsic polarization angle and in Faraday rotation, which occur both along the line of sight and across the telescope beam, cause depolarization of the radiation.

In this paper, it is more convenient to divide these depolarization mechanisms in depolarization along the line of sight and perpendicular to it, regardless of the physical process causing the depolarization. The latter is beam depolarization (Gray et al. 1999, Gaensler et al. 2001, Landecker et al. 2001, referred to as geometrical depolarization in Beck et al. 1999), and is two-dimensional, occurring in the telescope. Depolarization along the line of sight is depth depolarization (Landecker et al. 2001, Uyaniker and Landecker 2002, referred to as front-back depolarization by Gray et al. 1999), and is a one-dimensional addition of polarization vectors, assuming an infinitely narrow telescope beam.

### 3.3 The observations

In this study, we use two fields of observation, described in detail in Chapters 4 and 5. Some properties of the observations are described in Table 3.1. The observations were performed with the Westerbork Synthesis Radio Telescope (WSRT) at 8 frequencies near 350 MHz simultaneously, with a frequency bandwidth of 5 MHz. However, only 5 frequency bands, viz. those at 341, 349, 355, 360, and 375 MHz were usable due to radio interference. The regions are about 60 square degrees in size, both at positive Galactic latitudes, and mapped with a resolution (FWHM) of 5.0'×5.0' cosec δ. The regions were selected on the basis of polarization maps obtained in the WENSS survey at 325 MHz (Rengelink et al. 1997) because of their extraordinary small-scale structure in polarization, and were reobserved in multiple frequencies and higher sensitivity.

Standard data reduction techniques were used to obtain the Stokes parameters Q, U and I, from which polarized intensity $P = \sqrt{Q^2 + U^2}$ and polarization angle $\phi = 0.5\arctan (U/Q)$ were derived. $P$ and $\phi$ together form the complex polarization pseudo-vector $\mathbf{P} = P \exp(-2i\phi)$. The WSRT was used in mosaicking mode, i.e. the telescope...
repeatedly cycled through a grid of pointings a large number of times during one 12 hour observation, integrating 30 seconds per pointing each cycle. This yields a larger field of view with constant sensitivity, and lower off-axis instrumental polarization. See Chapter 2 for details on the data reduction.

Fig. 3.1 shows the polarized intensity distributions of the two regions. The structure in $P$ is uncorrelated with total intensity $I$. Rotation measures were derived from the linear relation $\phi = \phi_0 + RM\lambda^2$. Polarization angles are ambiguous on $\pm 180^\circ$, which has to be taken into account in determination of $RM$. The maximum $RM$ that can be obtained with minimum difference between angles at different frequencies is $RM_{max} = \phi_{max}/\Delta\lambda^2 = \pi/0.13 = 24$ rad m$^{-2}$. As the range in obtained $RM$ values is $|RM| \leq 15$ rad m$^{-2}$, ambiguities in polarization angle do not play any role in our observations. Furthermore, these low $RM$ values rule out any influence of bandwidth depolarization, which is the process of depolarization due to differential Faraday rotation across a (wide) frequency band. The linear $\phi(\lambda^2)$-relation can be destroyed by depolarization, which yields incorrect $RM$ values. Therefore, we only consider "reliably determined" $RM$ values, where "reliable" is defined as (1) the reduced $\chi^2$ of the linear $\phi(\lambda^2)$-relation $\chi^2_{red} < 2$, and (2) the polarized intensity averaged over frequency $\langle P \rangle > 20$ mJy/beam ($\sim 4 - 5\sigma$). $RM$ maps are given in Fig. 3.2, where filled circles denote positive $RM$s and open circles indicate negative $RM$s. The diameter of the symbol is proportional to the magnitude of $RM$. The $RM$ in the Auriga region (left plot) shows a clear gradient of about 1 rad m$^{-2}$ per degree in the direction of position angle $\theta = -20^\circ$ (N through E).

Histograms of the distributions of Stokes parameters $Q$ and $U$, total intensity $I$ and $RM$ are given in Fig. 3.3. In the $I$ map, point sources were subtracted down to 5 mJy/beam. For $Q$, $U$ and $I$, data from all five frequencies are shown in the same plot. The $RM$ plot of the Auriga region contains the histogram of all $RM$s as they are given in the $RM$ map in Fig. 3.2 (dashed line), as well as the histogram of $RM$ values where the best-fit linear gradient in $RM$ has been subtracted (solid line).

The histograms contain only statistical information on the two regions, and do not include the topology of the observed structure. The statistical information along these two separate lines of sight will be used in Section 3.10 to infer information on the Galactic magnetic field and correlation lengths in the warm ionized ISM. However, these two regions were selected because they show conspicuous structure in polarization, so they may exhibit more structure than an "average" part of the ISM away from the Galactic plane.

3.4 The effect of missing short spacings in aperture synthesis observations

In aperture synthesis observations, structure on large angular scales is not well represented because visibilities cannot be measured on baselines smaller than the diameter of the primary elements. For the Westerbork telescope, the shortest baseline is 36m, so that at 350 MHz, structure on angular scales larger than about a degree is not adequately measured. The proper way to correct for this undetectable large-scale structure is to observe the same region at the same frequencies with a single-dish telescope with
Figure 3.1: The observed regions in the constellation of Auriga (left) and of Horologium (right) in polarized intensity $P$ at 349 MHz at a resolution of 5.0'. $P$ is saturated at 95 mJy/beam (white), which corresponds to a polarized brightness temperature $T_{p,\text{sat}}$ of 12 K for Auriga, and 13.9 K for Horologium. The white boxes in the maps are regions of low polarization, that will be discussed in Section 3.6.

absolute intensity scaling and add these large-scale data to the interferometer data (see Stanimirovic (2002) for methods of data addition). However, for the WSRT at 350 MHz this is not possible, as there is no instrument of suitable size operating at these frequencies.

In the data reduction process of the WSRT, the lack of information on scales larger than about a degree is dealt with by setting the average value of measured intensities over the whole field to zero. For a strong source this will result in an image which has a bowl-like depression around the source, but for approximately uniform diffuse emission, it produces a more or less constant offset. In the case of polarimetry, this means that the average Stokes $Q$ and $U$ components are set to zero. So in each observed $Q$ and $U$ map there may be constant offsets $Q_0$ and $U_0$ that have to be added to the observed $Q$ and $U$ to obtain the real linearly polarized signal on the sky. Since the large mosaics are produced from several tens of pointings, each of which can have its own offsets, the offsets can vary over a mosaic.
The diffuse polarized radio background

Figure 3.2: Rotation measure maps of the observed regions in Auriga (left) and in Horologium (right). Rotation measures are denoted by white circles, where filled circles are positive RMs. The diameter of the symbol represents the magnitude of RM, where the scaling is given in rad m^{-1}. Only RMs for which \langle P \rangle \geq 5\sigma and reduced \chi^2 of the linear \phi(\lambda^2)-relation < 2 are shown, and only every second independent beam.

3.4.1 The effect of offsets on polarized intensity and rotation measure

The presence of offsets can create spurious small-scale structure in observed P, as sketched in Fig. 3.4. The left plot gives a simple one-dimensional example of a change in polarization angle, which causes small-scale structure in the distribution of Q_{true} and U_{true}, but not in P (center plot). The right plot shows the response of an interferometer: the average Q and U over the field are subtracted from the signal on the sky. P_{obs}, which is computed from Q_{obs} and U_{obs} does show apparent structure on small scales, although in reality it does not have that structure.

Furthermore, offsets will influence the RM determination, and will in general destroy the linear relation between polarization angle \phi and \lambda^2. Although in interferometer observations the Stokes Q and U emission can be separated in (observable) small-scale structure and (unobservable) large-scale structure, this is in general not true for the rotation measure, due to the complicated relation between intensity at different frequencies and RM.

A simple example of how offsets may distort RM is given in Fig. 3.5. Six plots in the (Q,U)-plane are drawn, each with five polarization pseudo-vectors. Each pseudo-vector denotes a certain wavelength, where the five wavelengths are equally spaced in \lambda^2, and are numbered according to increasing \lambda^2. The upper three plots give a hypothetical situation of three adjacent observed positions, where the pseudo-vectors denote the true polarization as it is on the sky. All three RMs are chosen to be positive, and the
Figure 3.3: Histograms of (from left to right) $Q$, $U$ and $I$ for 5 frequencies, and $RM$. $Q$, $U$ and $I$ data are oversampled, while only reliably determined $RM$s are included. Top plots are for the Auriga region, bottom plots for the Horologium region. In the solid line histogram of $RM$ in the Auriga region, the $RM$ gradient over the region is subtracted; the dashed line gives the histograms of the observed $RM$ including the gradient.

Figure 3.4: Example of how offsets can cause small-scale structure in $P$. Left: an example polarization angle distribution in one dimension. Center: small-scale structure in $Q$ (solid) and $U$ (dashed) corresponding to the change in polarization angle, while $P$ (dotted) remains constant. Right: the interferometer response to this distribution, where $P$ does show apparent structure.

value of $RM$ in the left plot is doubled and tripled in the central and right hand plot, respectively. Of these three plots, the $Q$ and $U$ values averaged over the 3 positions for each band separately were subtracted after which the polarization pseudo-vectors in the lower plots were obtained. Below these are plots of $\Delta \phi$ against the wavelength difference between two adjacent wavelengths $\Delta \lambda^2$, which gives the apparent $RM$. The
linear $\phi(\lambda^2)$-relation is thoroughly destroyed, the measured depolarization becomes highly non-linearly wavelength dependent, and the apparent $RM$ deviates from the true $RM$.

The fact that offsets can create structure in $P$ and prohibit reliable $RM$ determinations has to be given serious consideration in all interferometer observations with missing short spacings. We will estimate the importance of offsets in our observations in the next Sections. In Section 3.4.2 we discuss the case of a pure Faraday screen that is irradiated by a (constant) polarized background emission. Here, we will calculate analytically two limiting cases of broad and narrow $RM$ distribution, and then compute numerically the influence of offsets on $RM$ determinations. In Section 3.4.3 we expand this argument to treat a medium which contains both synchrotron-emitting and Faraday-rotating material.
3.4.2 Offsets for a uniformly polarized background propagating through a small-scale Faraday screen

First we consider the situation of a turbulent Faraday screen, i.e. a constant polarized background that undergoes Faraday rotation while propagating through a magnetized ionized medium. In this situation, small-scale structure in polarization angle is created by the Faraday rotation, while the polarized intensity remains unaltered. We can calculate the missing large-scale structure contribution for two limiting cases of broad and narrow width of the RM distribution $\sigma_{RM}$, assuming a uniform polarization background $P_0 = P_0 \exp(-2i\phi_0)$ where $\phi_0$ is the intrinsic polarization angle. Furthermore, we assume that the offsets can be approximated by a constant over the whole field of observation. The expected offsets can then be derived, depending on the RM distribution in the screen. We assume that the Faraday screen consists of cells with random RM, drawn from a Gaussian RM distribution of width $\sigma_{RM}$ and thus Faraday-rotates the background polarization angle on small scales. The offsets are then the normalized mean of $P_0$ weighted with the Gaussian RM distribution:

$$P_{\text{offsets}} = \frac{\int_{-\infty}^{\infty} P_0 \ e^{2i(\phi_0 + RM, \lambda^2)} e^{-RM^2/2\sigma_{RM}^2} \ dRM, \ \ (3.1)}{\int_{-\infty}^{\infty} e^{-RM^2/2\sigma_{RM}^2} \ dRM,}$$

This expression is independent of the angular length scale of the structure in RM, as long as the length scale is small enough with respect to the path length to have a Gaussian distribution of RMs. This yields the offsets $Q_0$ and $U_0$

$$Q_0 = P_0 e^{-2\sigma_{RM}^2 \lambda^4} \left[ \cos(2\phi_0) - \frac{1}{\sqrt{\pi}} \sin(2\phi_0) \int_0^{\sqrt{2\sigma_{RM}^2 \lambda^4}} e^{t^2} \ dt \right] \ (3.2)$$

$$U_0 = P_0 e^{-2\sigma_{RM}^2 \lambda^4} \left[ \sin(2\phi_0) - \frac{1}{\sqrt{\pi}} \cos(2\phi_0) \int_0^{\sqrt{2\sigma_{RM}^2 \lambda^4}} e^{t^2} \ dt \right] \ (3.3)$$

The offsets depend highly non-linearly upon the width of the random RM distribution $\sigma_{RM}$. The exact values can be easily calculated for two extremes:

a) The width of the $\phi$-distribution $\sigma_\phi = \sigma_{RM} \lambda^2 \geq \pi$, or large $\sigma_{RM}$:

$$e^{-2\sigma_{RM}^2 \lambda^4} \to 0 \quad \Rightarrow \quad Q_0 = U_0 = 0$$

The observed distribution of polarization angles is random, therefore $Q$ and $U$ are centered around zero and there is no undetected large-scale structure.

b) $\sigma_{RM}$ and accompanying $\sigma_\phi$ are so small that $\sigma_{RM} \lambda^2 \ll 1$:

$$e^{-2\sigma_{RM}^2 \lambda^4} \to 1 \quad \Rightarrow \quad \begin{cases} Q_0 = P_0 \cos(2\phi_0) \\ U_0 = P_0 \sin(2\phi_0) \end{cases}$$

Here the subtracted component is equal to the uniform component of the polarization vector. The observed polarized intensity is much lower than the true polarized intensity because of these large offsets.
A constant background rotation measure $RM_b$ can be incorporated by replacing $\phi_0$ in the equations by $\phi_0 + RM_b \chi^2$, but this does not change the results.

To study the effect of offsets on the observed $RM$ for intermediate values of $\sigma_{RM}$, we consider again a constant background polarization propagating through a random Faraday screen and estimate the difference between the $RM$ obtained in an absolutely calibrated single dish measurement, and the $RM$ observed with an interferometer. The Faraday screen is represented by a grid of $RM$s drawn randomly from a Gaussian distribution with width $\sigma_{RM}$. The high-pass filter applied to the interferometer is modeled by subtracting the mean values of $Q$ and $U$ over the field from the real values, to obtain the “observed” $Q_{obs}$ and $U_{obs}$ in 5 frequencies, just as in the Auriga and Horologium regions. The resultant apparent $RM$ from $Q_{obs}$ and $U_{obs}$ for 3 different values of $\sigma_{RM}$ are given in Fig. 3.6, as a function of the true $RM$ on the sky (that can be observed with a single dish).

The left plot of Fig. 3.6 shows the limiting case b) above, where the width of the $RM$ distribution is very narrow ($\sigma = 0.7 \text{ rad m}^{-2}$), which makes the average values of $Q$ and $U$ non-zero. Therefore the decrease in polarized intensity is large, and the output $RM$ deviates from the input $RM$. If $\sigma_{RM} \leq 1 \text{ rad m}^{-2}$, the decrease of polarized intensity is about 50% and the difference between $RM_{out}$ and $RM_{true}$ is of the order of $\sigma_{RM}$ but no larger, for frequencies around 350 MHz.

The central and right plot in Fig. 3.6 show the offsets if $\sigma_{RM} = 1.3 \text{ rad m}^{-2}$ and $\sigma_{RM} = 1.8 \text{ rad m}^{-2}$, respectively. The figure shows that, in the case of a Faraday screen irradiated by a polarized background, offsets only have to be dealt with if the width of the $RM$ distribution over the region is very small, typically $\sigma_{RM} \leq 1.8 \text{ rad m}^{-2}$.

### 3.4.3 The importance of offsets in the observations

If our observations were produced by a uniformly polarized background viewed through a pure Faraday-modulating screen, the observed width of the distribution of $RM$s (i.e. mostly $\sigma_{RM} \geq 1.8 \text{ rad m}^{-2}$) implies that offsets introduced would be negligible.
Auriga | Horologium
---|---
56.2° | 68.5°
55° | 67.3°
53.7° | 66°
52.5° | 64.8°
51.2° | 63.5°
50° |
48.7° |
90.6° | 94.6°
92.5° | 90.5°
93.4° |

Table 3.2: Width of RM distribution for each pointing in Auriga (left) and Horologium (right). The numbers within the boxes give $\sigma_{RM}$ in rad m$^{-2}$, and the coordinates denote the centers of the pointings.

![Graphs of RM distributions](image)

Figure 3.7: Example of RM distributions in separate pointings. The plots show the central row of pointings in the Auriga region. Dotted lines are Gaussian fits to the data, with the fitted $\sigma_{RM}$ are given above the plot.

in most pointing centers, as can be seen from Table 3.2. This table gives the values of $\sigma_{RM}$ in each pointing position in Auriga and Horologium in the boxes, while the coordinates outside the boxes give the pointing centers. Fig. 3.7 shows an example of the RM distributions in the central row of pointings in the Auriga field. In the majority of pointing centers $\sigma_{RM} \gtrsim 1.8$ rad m$^{-2}$, indicating negligible offsets. However, in Horologium, three pointings have $\sigma_{RM} \approx 1.6$ rad m$^{-2}$ and one pointing even $\sigma_{RM} \approx 1.3$ rad m$^{-2}$, which means that offsets could be present in these pointings of at most about 10% and 30% of the signal, respectively. It is not possible to compute these offsets so accurately that the correction for offsets can be made, so we have to take care in interpreting data from these pointings.

However, the observations show that the variation of observed polarization angle with wavelength does not perfectly follow the linear relation $\phi = \phi_0 + RM\lambda^2$, as one would expect for pure Faraday rotation (see Chapters 4 and 5). Offsets in $Q$ and/or $U$ can cause deviations in the linear $\phi(\lambda^2)$-relation, so we performed an additional test in the Auriga and Horologium regions to see if the observed non-linearities in the $\phi(\lambda^2)$-relation can be resolved by adding offsets.

Subfields of $\sim 1° \times 1°$ were selected from the data, centered around pointing centers. Large-scale $Q$ and/or $U$ components, independent for each of the five frequencies, were
added to the data to minimize the $\chi^2$ of the $\phi(\chi^2)$-relation. Offsets could be found in some subfields with a magnitude of the same order as $P$ in the subfield, which decreased the average $\chi^2_{red}$ by a factor of two. However, Fig. 3.8 shows the distribution of individual $\chi^2$ values per beam for one particular pointing for both Auriga (top) and Horologium (bottom). The left plots show $\chi^2$ values computed with offsets against $\chi^2$ without offsets. In the right plots, the histograms of $\chi^2$ without (solid line) and with (dotted line) offsets are given. In the Auriga pointing shown, offsets only cause a decrease in $\chi^2$ in 52% of the beams, although the average $\chi^2$ diminishes. In other pointings, this percentage ranges from 49% to 71%. Therefore, the computed offsets do not give a real improvement of the data, and cannot be considered real missing large-scale components. We acknowledge that this method has some caveats. First of all, it assumes that offsets are the only process distorting the linear $\phi(\chi^2)$-relation, while depolarization mechanisms can yield non-linearity too. In addition, it assumes that the offsets are constant over the probed subfield, which may not be true. Probing smaller subfields is no solution for this problem as the number of data points becomes too small with respect to the number of free parameters.

A third argument against dominant offsets in the data is the high quality of the determination of $RM$, i.e. a linear $\phi(\chi^2)$-relation with a low $\chi^2$. In $\sim 24\%$ of all data in the Auriga region and $\sim 31.3\%$ of the data in Horologium the reduced $\chi^2 < 2$ and average $P$ over all bands $> 20$ mJy/beam. However, of all pixels with $\langle P \rangle > 20$ mJy/beam, $\sim 70\%$ (in Auriga) and $\sim 62\%$ (in Horologium) has a $\chi^2_{red} < 2$. If offsets of the same order of the data would exist, $RM$s could not be so well-determined over such a large part of the fields.

Finally, models of a synchrotron-emitting and Faraday-rotating medium, presented in Section 3.9, do not show average $Q$ or $U$ values higher than about 10 mJy/beam.
From the large $\sigma_{RM}$, the good quality of the $RM$ determinations and minimizing $\chi^2$, we conclude that the presence of considerable missing large-scale structure is unlikely. In that case, the small-scale structure in polarized intensity must be due wholly to depolarization mechanisms. However, for a pure Faraday screen the only kind of depolarization that is possible is beam depolarization, because the observed values of $RM$ imply that bandwidth depolarization is not important, while depth depolarization requires that the rotating medium emits as well (see Section 3.9).

However, as shown in Section 3.6, beam depolarization — although its effect is clearly visible in the data — cannot explain all of the structure in polarized intensity. Therefore we are led to consider the not unrealistic situation in which we observe a polarized background that is modulated by a layer that both causes Faraday rotation, and emits synchrotron radiation. From the analysis in Section 3.7 it appears that the argument which limits the importance of offsets through the width of the distribution of observed $RMs$ applies equally to a pure Faraday screen and to a rotating and emitting screen.

The only way in which offsets could play a rôle is if there is a layer in front of the rotating and emitting screen which emits polarized radiation that is constant over the primary beam of our observations. A foreground-offset decreases the degree of polarization with a constant factor and can contribute a constant $RM$ component which cannot be derived from the data. However, a uniformly polarized foreground cannot influence the width of observed $RM$ distribution or induce small-scale depolarization. Judging from earlier single-dish data of $RM$ in the regions of the Auriga and the Horologium region (Bingham and Shakeshaft 1967, Spolstra 1984), we conclude that a possible undetected $RM$ component on scales $\gtrsim 1^\circ$, if present at all, must be very small.

### 3.5 Depolarization canals

A conspicuous feature in the observed polarized intensity is the presence of one-dimensional filament-like structures of low polarized intensity $P$. These so-called depolarization canals have been observed in many diffuse polarization observations (Wieringa et al. 1993, Duncan et al. 1999, Gray et al. 1999, Uyama et al. 1999, Gaensler et al. 2001). Two characteristics of the canals in our observations (and in others as far as we could judge from figures) are (1) the canals are one beam wide, and (2) the polarization angle changes across the canal by $90^\circ$. This typical behavior can be explained by two mechanisms:

1. The canals can denote a boundary between two regions of approximately constant polarization angle. A difference in polarization angle between the two regions of $\Delta \phi = (n + 1/2)180^\circ$ ($n = 0, 1, 2, \ldots$) will cause beam depolarization at the position of the boundary, i.e. depolarization due to varying polarization vectors across the telescope beam (Haverkorn et al. 2000, Chapter 6). This boundary is by definition one beam wide. We expect boundaries between regions of constant polarization angle for every value of $\Delta \phi$, but the only $\Delta \phi$ that is clearly observable in $P$ maps is $\Delta \phi = (n + 1/2)180^\circ$. This means that there is no fortuitous
The diffuse polarized radio background

abundance of \( \Delta \phi = (n + 1/2) 180^\circ \), but it is the only \( \Delta \phi \) value which is visible in
maps of polarized intensity.

2. A medium containing uniform magnetic field and electron density depolarizes
polarized radiation by means of differential Faraday rotation, see Section 3.2. In
this case, the observed polarized intensity is

\[
P = P_0 \left| \frac{\sin(2RM \lambda^2)}{2RM \lambda^2} \right| \tag{3.4}
\]

(Burn 1966, Sokoloff et al. 1998), where \( P_0 \) is the polarized intensity observed
at \( \lambda = 0 \). A linear gradient in RM would produce long narrow depolarization
canals at a certain value \( RM \), where \( 2RM, \lambda^2 = n\pi \). Across every null in the
sinc-function, the polarization angle changes by \( 90^\circ \).

We will attempt to estimate the relevance of both explanations to our observations
below.

3.5.1 Frequency dependence of canals

If the canals are formed by beam depolarization, there are two extreme possibilities for
the origin of the angle change \( \Delta \phi = (n + 1/2) 180^\circ \): either the \( RM \) of the ISM changes
across the canal by an amount of \( (n + 1/2) 180^\circ / \lambda^2 \), or the \( RM \) of the ISM does not
change and there is an intrinsic angle difference of \( \Delta \phi_0 = (n + 1/2) 180^\circ \). These two
extremes cannot be distinguished from observations at a single frequency. However, on
the basis of the first hypothesis, one would expect the canals to change in intensity with
frequency, while they should be identical at different frequencies if the \( RM \) does not
change. On the other hand, if the canals are caused by differential Faraday rotation,
the polarized intensity in a canal should change with frequency according to Eq. (3.4).

We can test the frequency dependence of the depth of the canals as follows. Canals
are defined as sets of “canal-pixels”. A pixel is defined as a “canal-pixel” if the polarized
intensity is low \( P < 2 \) times rms noise\) and \( P \) on diametrically opposed sides of
that pixel, one beam away, is high \( P > 5 \) times rms noise\). The high-\( P \) pixels
surrounding the canal-pixel can be oriented horizontally, vertically or diagonally. No
further assumptions regarding the length of canals are made; therefore a single pixel
with low \( P \) that is not part of a canal but is surrounded by high \( P \) pixels, is also
included in the definition of a canal-pixel. Five sets of canal-pixels are evaluated, for
each frequency separately, and the average \( P \) in each set of canal-pixels is computed for
all five bands. We plot this against \( \lambda^2 \) to probe the change in the depth of the canals
over frequency. For each of the two regions we show the average \( P \) for the canal-pixels
in Fig. 3.9, defined in each of the 5 frequency bands, i.e for 341 (left), 349, 355, 360
and 375 MHz (right). The dashed lines show the predictions of \( P(\lambda^2) \) for canals that
are caused by beam depolarization, and are due to a change in \( RM \) (with \( \Delta RM = 2.1, 6.3 \)
and 10.5 rad m\(^{-2}\) for the three lines, respectively). Note that for canals that are
caused by beam depolarization due to an intrinsic change in angle, with constant \( RM \),
\( P(\lambda^2) \) is constant. The dotted lines in Fig. 3.9 denote the prediction of the polarization
angle if the canal is caused by differential Faraday rotation.
Figure 3.9: Plots of the average $P$ of a set of pixels contained in canals against wavelength, in the Auriga region (top) and Horologium region (bottom). In each plot the canals were defined in an other frequency band, at 341, 349, 355, 360, and 375 MHz from left to right. The behavior of the $P$ distribution if all canals were caused by beam depolarization due to a change in $RM$ is shown in dashed lines, from bottom to top for $\Delta RM \approx 2.1, 6.3, 10.5$ rad m$^{-2}$ (i.e. $\Delta \phi = 90^\circ, 270^\circ, 450^\circ$). The behavior of $P$ if the canals were caused by differential Faraday rotation are shown dotted for $2RM\lambda^2 = n\pi/2$ for $n = 1, 3, 5$.

In both regions, the wavelength in which the canals are selected has the lowest average $P$ for each frequency, with $P$ increasing with $\Delta \lambda^2$. This rules out the possibility that the canals are caused by a change in intrinsic angle, confirming the conclusion from the non-detection of $I$ that the background polarized intensity is smooth. As the predictions in dashed and dotted lines are arbitrarily normalized, they can be shifted up and down. However, the shape of the predicted lines is typical of the responsible depolarization process. Judging solely from the shape of the lines, the explanation of differential Faraday rotation seems to make a less bad fit, although none of the two models accurately predict the observations.

3.5.2 $RM$ and $\Delta RM$ values at canals

Canals due to beam depolarization are caused by a specific change in $RM$ across a canal $\Delta RM_c = (n + 1/2)\pi/\lambda^2$. On the other hand, canals caused by differential Faraday rotation are determined by a specific absolute $RM_c = n\pi/(2\lambda^2)$. With the sets of canal-pixels defined in the previous subsection, we define $\Delta RM_c = RM_1 - RM_2$, where 1 and 2 are high-$P$ pixels on opposite sides of the canal. Then at the canal-pixel, $RM_c = (RM_1 + RM_2)/2$. The observed distributions of $\Delta RM_c$ and $RM_c$ are shown in Fig. 3.10. There are indications that both $\Delta RM_c$ and $RM_c$ show peaks in
Figure 3.10: Distribution of \( \Delta R.M. \) across a canal and the absolute \( R.M. \) at a canal (as estimated from its neighbors), where the canal is defined at frequencies 341, 349, 355, 360, and 375 MHz from left to right. Top panels show the Auriga region and bottom panels the Horologium region. Only \( R.M. \) values where \( \chi^2_{red} < 2 \) and \( P > 30 \) mJy/beam are used. In the \( \Delta R.M. \) plots, dotted vertical lines are \( \Delta R.M. \) values where \( \Delta \phi \) across a canal would be ±90°, ±270° or ±540°. In the \( R.M. \) plots, the dotted vertical lines are values where \( 2 R.M., \lambda^2 = n\pi/2 \).
the distribution at the values needed to create canals, although clearly the observations show no perfect agreement with any of the two methods. Note that canals with angle changes $\Delta \phi \leq 90^\circ$ have slightly different accompanying $\Delta RMs$ or $RMs$ and therefore broaden the peaks. Noise in $RM$ has the same effect.

3.5.3 Position shift of the canals with frequency

Differential Faraday rotation depolarizes all lines of sight which have a certain $RM$ value $RM_c$. This means that depolarization is not necessarily in the form of narrow one-dimensional canals, but could also be caused in two dimensions, e.g. by a cloudlet causing $RM = RM_c$. The easiest way to explain one-dimensional canals in this approach is assuming a gradient in $RM$, so that the $RM$ stays constant over a certain length perpendicular to the gradient, and a one-dimensional canal is formed. But at 375 MHz, the canal will form at beams where $RM_c = 2.46 \text{ rad m}^{-2}$ while it forms where $RM_c = 2.02 \text{ rad m}^{-2}$ at 341 MHz. If we assume typical gradients of 1 rad m$^{-2}$ per degree, similar to the large-scale gradient observed in the Auriga region, this indicates that the canal should move with position over $\sim 5$ beams from 341 MHz to 375 MHz. Instead, canals move at maximum 3 pixels from 341 to 375 MHz, which is about 0.5 beam. A gradient in $RM$ would have to be larger than $\sim 13 \text{ rad m}^{-2}$ per degree to position the canals in the 5 frequencies within half a beam from each other. Such a gradient is certainly possible locally, although this high gradient would have to extend over a large part of the field in order to explain the long and straight canals. Furthermore, if such large gradients are present in the medium, we would expect lower gradients as well. These lower gradients would give canals that shift position with frequency significantly, which is not observed.

3.5.4 Shape of the canals

Because the data is oversampled by a factor of six, it is possible to determine the distribution of $P$ in a cross-cut through the canals.

If the canals are caused by beam depolarization, we can estimate the gradient of the polarization angle within the beam, because a large gradient causes a deeper canal than does a smaller gradient, as illustrated in Fig. 3.11. Here a one-dimensional example is
given of a change in polarization angle (left plot) and the corresponding change in \( P \), after convolution with the telescope beam. The narrowest \( P \) profile is achieved when the change in angle is on a length scale smaller than about 20% of the beam, or about one pixel.

Such narrow \( P \) profiles are indeed observed across canals, as is shown in Fig. 3.12. In this figure the top plots give a one-dimensional cross-cut of \( P \) across a canal against position, for all frequencies. The frequencies in which the canals were defined were 341 MHz, 355 MHz and 349 MHz respectively. The bottom plots give only the \( P \) distribution across the canal at the frequency at which it was defined (solid line). Superimposed in dashed lines is the \( P \) distribution of the model of Fig. 3.11 for the steepest angle change convolved with the synthesized beam. Less steep angle changes give less steep \( P \) profiles and worse fits to the data.

In this case, changes in polarization angle of at least 90° must exist on angular scales smaller than an arcminute in about 10 - 20% of all canals, which are mostly accompanied by an abrupt change in the \( R M \) computed from the observed \( Q \) and \( U \). As \( R M \) is the integral of magnetic field and electron density over path length, it is not easy to understand an abrupt change of \( \Delta R M / R M \approx 100\% \) over an arcminute, i.e. about a pixel.

On the other hand, an interpretation of these “sharp” canals in terms of differential Faraday dispersion has its problems as well. First, the canals would have to be much more closely spaced than observed. Secondly, only a medium uniform in both magnetic field and electron density can produce such narrow and deep canals. Sokoloff et al. (1998) show that an exponential asymmetric slab causes non-zero minima for the canals, which even disappear completely in a turbulent medium.

### 3.5.5 Beam depolarization or differential Faraday rotation?

We conclude that at this stage we cannot distinguish decisively between beam depolarization and differential Faraday rotation as an explanation for canals, although we believe that the latter explanation is more difficult to reconcile with the data.

If beam depolarization is responsible for the canals, abrupt \( R M \) changes have to be present in the medium. Because \( R M \) is an integral along the line of sight, it is difficult to see what physical process would be responsible for this. However, numerical models by Sokoloff et al. (1998) show that abrupt changes in apparent \( R M \) are possible for a turbulent medium. Furthermore, beam depolarization can create abrupt \( R M \) changes across a beam as well (Chapter 9). Nevertheless, Fig. 3.10 shows that this certainly is not the whole explanation.

If differential Faraday rotation creates the canals, then it is hard to understand why all canals are exactly one beam wide, and why we do not observe any variation in the position of the canals with frequency. Furthermore, the observation of canals almost down to \( P = 0 \) indicates a very uniform medium in both magnetic field and electron density. Small-scale structure in observed \( R M \) indicates that this is not the case. For a non-uniform medium, canals become less pronounced, or even disappear.
Figure 3.12: Upper plots: examples of observed one-dimensional $P$ distributions in the Horologium region for five frequencies, where the deepest canal is observed at 341, 355 and 349 MHz respectively. Lower plots: the same $P$ distribution for the deepest canal as above (solid) and the best fit according to the model of Fig. 3.11. The $P$ profile is so steep that the change in angle that causes the canal must be on scales of an arcminute or smaller.

3.6 Beam depolarization

Beam depolarization can be recognized because of its characteristic angular scale, viz. that of the beam. The magnitude of the effect depends strongly on the variation of polarization angle within the beam. The latter cannot be measured, but a lower limit to the amount of depolarization can be derived from the gradient of polarization angle on scales that are slightly larger than the size of the beam.

3.6.1 Beam depolarization at low signal-to-noise ratios

In the regions of low polarized intensity $P$, a mottled structure in $P$ is visible in both the Auriga and the Horologium regions on scales of a beam width. This is a consequence of beam depolarization acting on low signal-to-noise data. Assuming a random polarization angle distribution, convolution with a Gaussian telescope beam results in a pattern as shown in Fig. 3.13. The shape of the distribution of $P$ is independent of the angular scale on which the polarization angle changes significantly, expressed as a fraction of the size of the beam; it is always a Rayleigh distribution, as is expected for $P = \sqrt{Q^2 + U^2}$ from independent Gaussian $Q$ and $U$ distributions.

We selected regions of low polarization in both Auriga and Horologium fields, which
are indicated by the white boxes in Fig. 3.1. Estimates of the degree of polarization $p = P/I$ are based on the measured values of $P$, and on estimates of $I$ derived from the radio survey at 408 MHz by Haslam et al. (1981, 1982), see Section 3.8.1. This yields degrees of polarization between 5% and 7% for the regions of low polarization, whereas the average degree of polarization is $p = 11 - 12\%$ for both regions. (The degree of polarization due to noise in $Q$ and $U$ is less than 1.5%.)

If we assume a maximum degree of polarization of $\sim 70\%$ (Burn 1966) before vector averaging within the beam, we find that we need structure on scales smaller than 1/3 of the beam to produce the observed degree of polarization. Clearly, if the degree of polarization before vector averaging is less than 70%, less structure within the beam is required.

From the appearance of Fig. 3.13, we conclude that the mottled pattern visible in regions of low polarization in both the Auriga and Horologium regions is caused by beam depolarization of regions with very little coherent large-scale structure (i.e. over many beams) in polarization angle. If beam depolarization is the dominant depolarizing mechanism, the angle structure within the beam must be present on scales at least three times smaller than the beam to obtain the observed degree of polarization. If the sub-beam structure is not random but correlated, structure on even smaller scales is needed. Structure on sub-arcminute scales is found in electron density in the nearby ISM ($\lesssim 1$ kpc) from pulsar studies (Armstrong et al. 1995), and in $\Delta$Ms measured from extragalactic sources, which includes the entire path length through the medium (Clegg et al. 1992). However, random structure on scales much smaller than the beam width is unlikely to dominate in these observations, because coherent structure over many beams is observed as well. Therefore, other depolarization mechanisms most likely contribute significantly. The obvious candidate is depth depolarization which we will discuss below, in Section 3.7.
3.6.2 Beam depolarization due to a constant angle gradient

A constant gradient in polarization angle over the beam causes constant beam depolarization. For a resolved gradient (Sokoloff et al. 1998):

$$p = e^{-2 \Delta R M^2 \lambda^4} = e^{-2 \Delta \phi^2}$$

where $\Delta R M$ is the gradient in $R M$ in rad m$^{-2}$ per beam. At $\sim 350$ MHz, beam depolarization due to an angle gradient starts to become important with angle gradients of ten degrees to several tens of degrees per beam. These angle changes are common in both regions. Fig. 3.14 shows a part of the Auriga region where $P$ is relatively high. The grey scale is $P$ at 349 MHz, oversampled by a factor 5, superimposed are polarization vectors of independent beams. Angle changes per beam of $\Delta \phi = 20^\circ$ are common, causing a depolarization down to 78% of the original $P$ according to Eq. (3.5). Changes of polarization angle of 30$^\circ$ over a beam even cause a reduction to 57% of the polarized intensity. So, beam depolarization due to a smooth gradient in angle is very effective in reducing the degree of polarization at positions with a significant angle gradient on beam scales, although its contribution is hard to estimate quantitatively from the observations.

3.6.3 The importance of beam depolarization

An estimate of the importance of beam depolarization in our data can be obtained by assuming that there is no structure in the polarization angle within a beam that cannot be predicted from the neighboring beams. In this case, one can estimate the structure in polarization angle across the beam from the values of angles in the neighboring beams. This method overestimates the contribution of beam depolarization when the
polarization angle has a gradient on scales larger than a beam, because the method averages angles over a larger area than one beam. However, it underestimates the beam depolarization if there is additional structure in angle within the beam that cannot be predicted from neighboring beams. For each beam, we estimate the expected amount of beam depolarization from averaging the $Q$ and $U$ values in directly adjacent beams and we correct the observed polarized intensity in the beam for the amount of depolarization estimated in that way. The resulting corrected distribution of polarized intensity for two frequencies 349 MHz and 375 MHz is shown in Fig. 3.15 in the left-hand panel, with the observed $P$ in the right-hand panel for comparison.

In the maps that are corrected for beam depolarization, the almost complete disappearance of the canals is striking. Only in the central pixel of the canals the depolarization is not always correctly accounted for, because $P$ is so low there that the angle determinations become very noisy. This effect produces overestimates and underestimates of depolarization, which are still visible in the corrected data as little white speckles and weak, one-pixel ($\sim 1/6$ beam) wide canals. Regions of high $P$ in the corrected maps have higher $P$ than in the non-corrected maps because here the beam depolarization is overestimated. This is because in the model, polarization angles are averaged over a length of three beams instead of over one beam. Therefore, structure in polarization angle that is averaged over a length of three beams causes a larger depolarization than if averaged over one beam only.

However, at high $P$, the large-scale structure is very similar in the corrected and uncorrected maps. This illustrates the fact that beam depolarization operates only on the angular scale of the beam, but not on larger scales. So except for the canals, beam depolarization is not very important, and the large-scale structure in $P$ must be due to other effects. In the following sections we discuss the primary cause of the large-scale structure in $P$, namely depth depolarization.

3.7 Depth depolarization

The absence of correlated structure in $I$ appears to be a general feature: both regions discussed here confirm it, as well as other observations made with the WSRT at wavelengths around 90 cm (e.g. Katgert and de Bruyn 1999, and Chapter 7). The lack of corresponding structure in total intensity $I$ suggests that Faraday rotation is the main process responsible for the observed structure in polarization. However, the structure in $P$ cannot be produced by this Faraday modulation. As shown in Section 3.4, a missing large-scale structure component cannot be strong enough to create all the structure in $P$, and neither can beam depolarization (except for the canals), as shown in Section 3.6.

We are therefore led to consider the more realistic situation in which the medium that produces the Faraday rotation is mixed with a medium that emits synchrotron radiation, which produces structure in $P$ through depth depolarization. The low level of small-scale structure in $I$ in all of these observations implies that the total intensity of the synchrotron emission must be very uniform. However, many observations suggest that the magnetic field, and therefore the synchrotron emissivity of the thin disk, is far from uniform. Therefore, the smoothness of the synchrotron total intensity cannot be
Figure 3.15: Left: polarized intensity in Auriga at 349 MHz (top) and 375 MHz (bottom) corrected for beam depolarization from the observed angle distribution in directly adjacent beams (for details see text). Right: observed polarized intensity at 349 MHz (top) and 375 MHz (bottom). Depicted in all plots is the square root of $P$ to decrease the dynamical range of the plots. The display scale saturates at $P = 90$ mJy/beam.
due to homogeneous synchrotron emission. Instead, the number of turbulent cells along
the line of sight must be so large that the spatial variation in synchrotron emissivity
is averaged out.

The apparent contradiction between uniform total intensity $I$ and highly-structured
polarized intensity $P$ is resolved by the fact that in the second quadrant, where all
our observations are done, the uniform component of the Galactic magnetic field is
believed to be perpendicular to the line of sight (e.g. review by Beck, 2001). This
means that $B_{\perp}$ has a large uniform component, so that emitted synchrotron radiation
is mostly uniform in total intensity $I$. On the other hand, $B_{\parallel}$ is dominated by the
random, small-scale component of the field, which produces small-scale structure in
rotation measure.

We shall next discuss the ingredients of a model for a Galactic disk which contains
a mixture of rotating and emitting material, and a thick disk or halo providing a
constant polarized background. Section 3.8 briefly describes the components of the
ISM relevant to the model. In Section 3.9 we will describe the model in some detail,
and how observational constraints can be used to derive estimates for parameters like
the strengths of the perpendicular and parallel components of the regular magnetic
field, and the strength of the small-scale random magnetic field. In Section 3.10 we
apply the model to our observations.

3.8 Relevant components of the ISM

3.8.1 Cosmic rays and thermal gas

The synchrotron intensity is the integrated non-thermal emission along the line of
sight, and it depends on the relativistic electron density $n_{e, r, d}$ and magnetic field
perpendicular to the line of sight $B_{\perp}$. Beuermann et al. (1985) have modeled the Galactic
synchrotron emissivity $\varepsilon$ from the continuum survey by Haslam et al. (1981, 1982) at
408 MHz. They incorporate spiral structure in the synchrotron radiation and find two
distributions of emission: a galactocentric thick and thin disk, with half equivalent
widths of $h_{e, b} = 1800$ pc and $h_{e, n} = 180$ pc at the radius of the Sun, respectively
(assuming a galactocentric radius of the Sun $R_\odot = 10$ kpc) and increasing outwards.
The half equivalent widths are defined as $\varepsilon(z) = \varepsilon_0(0) \sech^b(z/z_0) + \varepsilon_n(0) \sech^n(z/z_0)$,
where $z_0$ is the half equivalent width of the stellar disk, and the exponents $b$ and $n$
together with $z_0$ define the half equivalent widths of the two disks. The thick disk
provides about 90% of the synchrotron total power. Beuermann et al. only use data
in the first and fourth quadrant to set the parameters in their model, and infer from
the model the distribution of Galactic synchrotron emission in the second and third
quadrant.

The major part of the Faraday rotation is caused by the warm ionized medium,
because of its high thermal electron density, contained in a medium called the Reynolds
layer (Reynolds 1989, Reynolds 1991). The layer has a height of about 1 kpc, a
temperature $T \approx 8000$ K, and a thermal electron density concentrated in clumps of
$n_{e, th} \approx 0.08$ cm$^{-3}$ with a filling factor of about 20%.

These two distributions divide the ISM in three regimes, as sketched in Table 3.3,
where the base is the Galactic plane and $z$ increases upwards. The first domain coi-
Table 3.3: Three schematic domains with different characteristics in the Galactic ISM and halo. The Galactic plane is at the bottom, and z increases upward.

<table>
<thead>
<tr>
<th>Domain</th>
<th>Characteristic</th>
<th>Height</th>
<th>Contains</th>
<th>Typical Cell Size</th>
</tr>
</thead>
<tbody>
<tr>
<td>III</td>
<td>Thick disk</td>
<td>1800 pc</td>
<td>$n_{rd}, B$</td>
<td>100 – 1000 pc</td>
</tr>
<tr>
<td>II</td>
<td>Reynolds layer</td>
<td>1000 pc</td>
<td>$n_{rel}, B, n_{th}$</td>
<td>100 – 1000 pc</td>
</tr>
<tr>
<td>I</td>
<td>Thin disk</td>
<td>180 pc</td>
<td>$n_{rel}, B, n_{th}$</td>
<td>10 – 100 pc</td>
</tr>
</tbody>
</table>

Galactic plane

cides with the thin synchrotron disk. This also coincides with the stellar disk (~ 200-300 pc), the thin HI disk (~ 200 pc, Dickey and Lockman 1990), and the disk of classical HII regions (~ 60 pc). The second domain is the layer above the thin disk which contains the Reynolds layer. So in domains I and II, relativistic and thermal electrons are mixed, causing depolarization effects. The third domain is the highest layer, in which the thermal electron density is negligible, so that there is no Faraday rotation, but relativistic electrons are present that emit synchrotron radiation. Note that in the direction of our observations, the scale height parameters are highly uncertain, so that in practice it is uncertain whether there is a layer without thermal gas.

The hot ionized component of the ISM can contribute to the Faraday rotation as well. By Faraday rotation alone, it is impossible to distinguish between warm and hot gas. However, the hot gas does not contribute to the Hα emission, so that the warm and hot gas can be separated with Hα observations.

3.8.2 Regular and random Galactic magnetic field

We decompose the Galactic magnetic field in a regular large-scale component and a random component $B = B_{reg} + B_{ran}$.

Estimates of the ratio of random to regular magnetic field strengths $B_{ran}/B_{reg}$ seem to depend on the method used. Magnetic field determinations using RMs from extragalactic sources yield $B_{ran}/B_{reg} \approx 0.5 – 1$ (Jokipii and Lerche 1969, Clegg et al. 1992), while pulsar RMs indicate that $B_{ran}/B_{reg} \approx 3 – 4$ (Rand and Kulkarni 1989, Ohno and Shibata 1993). From diffuse polarization measurements, Spoelstra (1984) estimates for the ratio between random and regular magnetic field strengths $B_{ran}/B_{reg} \approx 1 - 3$, in agreement with Phillips et al. (1981), who find from the diffuse synchrotron background that $B_{ran}/B_{reg} \geq 1$. Heiles (1996) estimates a average from several studies as $B_{ran}/B_{reg} \approx 2$.

Structure in the rotation measure is estimated to be present on scales from 0.1 to 100 pc from RM observations of extragalactic point sources (Clegg et al. 1992, Minter
and Spangler 1996), whereas pulsar rotation measures and dispersion measures give cell sizes of 10 to 100 pc (Ohno and Shibata 1993), and starlight polarization measurements even yield cell sizes of a kpc (Jones et al. 1992).

Field strengths and structure in the Galactic halo, i.e., in the gas above the thin synchrotron disk at \( h \gtrsim 200 \) pc, can be estimated from observations of halos of external galaxies. In observations of synchrotron emission in halos of edge-on galaxies, the degree of polarization mostly increases with distance from the plane, suggesting a decreasing irregular magnetic field component for increasing distance to the plane of the galaxy. Structure in the thick disk or halo varies on much larger scales than in the thin disk, viz. on scales of about 100 – 1000 pc (e.g., Dunke et al. 1995).

### 3.8.3 The Local Bubble and Local Interstellar Cloud

The Sun is embedded in the Local Bubble (LB), a cavity of hot \(( T \approx 10^6 \) K) gas with thermal electron density \( n_{e,th} \approx 4 \times 10^{-3} \) cm\(^{-3}\) (Snowden et al. 1990). The path length through the Local Bubble varies from \( \sim 100 \) to 300 pc, and is outlined in Fig. 6 of Snowden et al. (ibid.) in different projections. In the direction of the Auriga and Horologium regions, the path length through the Local Bubble is not more than 100 pc. So a \( RM \) contribution from the Local Bubble would be \( R_{MB} \approx 0.34 B_{||} \) rad m\(^{-2}\), where \( B_{||} \) is in \( \mu G \). Leroy (1999) argues that the magnetic field in the LB does not follow the large scale field but has structure on smaller scales, which results in a smaller \( B_{||} \), so in a smaller \( RM \). So the Local Bubble will contribute to the observed \( RM \), but likely not more than \( \sim 1 \) rad m\(^{-2}\).

The Local Bubble contains cloud complexes of warm ISM. The Sun is just inside one of these, the Local Interstellar Cloud (LIC), which extends to a distance of \( l \approx 5.5 \) pc in the second quadrant. The electron density in the LIC is \( n_e \approx 0.04 – 0.15 \) cm\(^{-3}\) (Linsky et al. 2000). So \( R_{MLC} \approx 0.45 B_{||} \) rad m\(^{-2}\) \(( B_{||} \) in \( \mu G \)), contributing a small amount to the \( RM \) if there is a considerable magnetic field component along the line of sight.

### 3.9 A model of a Faraday-rotating and synchrotron-emitting layer

In this section we describe a simple model of a thin Galactic disk containing cosmic rays, magnetic fields and thermal electrons (layer I in Table 3.3), irradiated by a uniform polarized background (from layers II and III in Table 3.3). We calculate the total intensity, Stokes \( Q \) and \( U \), and the implied rotation measure, for various assumptions about the structure of the layer. In Section 3.10, we will compare these results with the observations.

#### 3.9.1 Outline of the model

In the thin disk, structure in the warm gas and in the magnetic field exists, which we describe with a model with cells of a fixed size. In this model, only domain I defined in Table 3.3 has structure, while domains II and III are assumed constant. Fig. 3.16 gives a sketch of the model and its parameters; in addition to the cell size \( d \) these are \( h \), the
vertical height of the layer, and the synchrotron emissivity \( I_e \) in each cell. The warm ionized medium has a filling factor \( f \); this is accounted for in the model by setting the electron density to an assumed value \( n_e \) in a fraction \( f \) of the cells along the line of sight, which are randomly chosen. In the remaining fraction \((1 - f)\) of cells, \( n_e \) is set to zero. This means that we have made the simplifying assumption that the hot gas does not contribute significantly to the rotation measure.

The magnetic field in the thin turbulent disk consists of a random \( B_{ran} \) and regular \( B_{reg} \) component. As it is not very well known how the regular and random components of the magnetic fields in the cold, warm and hot phases of the ISM are related, we consider two extreme cases:

- **Model A**: both the random and the regular component of the magnetic field do not change throughout the ISM in the thin disk, regardless of the phase of the ISM.

- **Model B**: the random component of the magnetic field only exists in the turbulent warm ISM. In the cold and hot ISM, the regular magnetic field component predominates. For equilibrium reasons, the total magnetic field energy density is taken to be the same in all phases.

In each cell, an amount of synchrotron radiation \( I_e \propto B_e^2 \) is emitted. This emission, and the emission from each cell further down the line of sight and from the background passing through the cell, is Faraday rotated by an amount \( \phi_{Fr} \) (in radians). So in each
cell:

\[ I_c = \frac{C}{N} \left( (B_{ran} \sin \alpha)^2 + B_{reg,\perp}^2 \right) \]

\[ \phi_{Pr} = RMX^2 = 0.81 n_c (B_{ran} \cos \alpha + B_{reg,\parallel}) \, d\lambda^2 \]

\[ P_c = 0.7 I_c \]

where \( B_{ran} \) is the constant strength of the random magnetic field in \( \mu G \), \( \alpha \) its random angle with the line of sight, \( C \) a proportionality constant and \( N \) the number of cells along the line of sight. \( C \) is divided by \( N \) to ensure that the total emission from the layer \( \approx CN \) is comparable for different cell sizes. The polarized emission in each cell \( P_c \) equals the maximum polarization of synchrotron radiation \( I_c \) generated in a cell, which is related to the power spectral index \( \gamma \) of the electron energy distribution as \( p(\gamma) = (3\gamma + 3)/(3\gamma + 7) \) (Burn 1966). For \( \gamma \) around 2.7, the maximum polarization is \( \sim 70\% \) of the total intensity \( I \). The polarization angle of the polarized emission generated in each cell \( \phi_{in} \) is taken to be the position angle of the perpendicular magnetic field. The position angle of the random magnetic field component \( \phi_r \) is random, and that for the regular component is chosen in the direction of Galactic longitude. Therefore the polarized intensity emerging from a cell is

\[ P_c = 0.7I_c e^{-2i(\phi_{Pr} + \phi_{in})} + 0.7I_k e^{-2i(\phi_{Pr} + \phi_{in})} \]  \( (3.6) \)

for a cell that is irradiated with polarized intensity \( I_k \) and polarization angle \( \phi_k \).

The thick synchrotron disk (layers II and III in Table 3.3) serves as a background to the detailed model of layer I (the thin disk). The effect of the thick disk is modeled as follows. We assume that the structure in the thick disk is on such large scales that we can approximate the background as constant, and that it is a uniform emitter of synchrotron radiation. The thick disk produces a constant polarized intensity \( P_k \) as input to layer I, with uniform polarization angle. This polarized intensity has two components: in layer III (where the density of thermal electrons is assumed to be zero) there is no depolarization because the magnetic field is uniform. In layer II, depth depolarization in a uniform layer (equivalent to differential Faraday rotation) will decrease the degree of polarization. Since it is not the purpose of our model to give a detailed description of layers II and III, we used a single relation to describe the effect of the thick disk, viz. \( P_k = 0.7 \eta_k I_k \), with \( \eta_k \) a factor \( 0 \leq \eta_k \leq 1 \) describing the amount of uniform depolarization in the layers above the modeled thin disk.

A given line of sight through the model grid (to be identified with the direction of one of our fields, and corresponding to a particular Galactic latitude) is simulated many times, by independently filling the cells that contain the warm ISM, and by redrawing the angle that the random component of the magnetic field makes with the line of sight. An ensemble of such realizations, for which we derive the distributions of \( I, Q, U \) and \( RM \), simulates the many lines of sight for which we obtain the same information in one of our observational fields. However, only the statistical information of these distributions is included in the model, not the discrete structure.

Beam depolarization is not included explicitly in the models. However, all allowed cell sizes have an angular scale at the far edge of the layer larger than the resolution. Therefore, significant additional depolarization due to beam depolarization is not expected.
### Input parameters with fixed values

<table>
<thead>
<tr>
<th>Parameter</th>
<th>Value</th>
</tr>
</thead>
<tbody>
<tr>
<td>( n_c )</td>
<td>thermal electron density in cells</td>
</tr>
<tr>
<td>( f )</td>
<td>filling factor of the warm ISM</td>
</tr>
<tr>
<td>( h )</td>
<td>height of the layer with cells</td>
</tr>
<tr>
<td>( \phi_0 )</td>
<td>intrinsic polarization angle of the background</td>
</tr>
<tr>
<td>( \phi_s )</td>
<td>position angle of random ( B )-field</td>
</tr>
<tr>
<td>( \alpha )</td>
<td>angle between random ( B )-field and line of sight</td>
</tr>
<tr>
<td>( I_0 )</td>
<td>total intensity</td>
</tr>
<tr>
<td></td>
<td>0.08 cm(^{-3}) (Reynolds 1991)</td>
</tr>
<tr>
<td></td>
<td>20% (Reynolds 1991)</td>
</tr>
<tr>
<td></td>
<td>180 pc (Beuermann et al. 1985)</td>
</tr>
<tr>
<td></td>
<td>arbitrary: 0(^\circ) chosen</td>
</tr>
<tr>
<td></td>
<td>random per cell</td>
</tr>
<tr>
<td></td>
<td>random per cell</td>
</tr>
<tr>
<td></td>
<td>Auriga: 34 K</td>
</tr>
<tr>
<td></td>
<td>Horologium: 47 K</td>
</tr>
<tr>
<td></td>
<td>from Haslam et al. (1982)</td>
</tr>
</tbody>
</table>

### Free input parameters

- \( d \): cell size
- \( C \): proportionality constant between synchrotron emission and \( B_1^2 \)

### Constraints determined from the observations

- \( R \bar{M}_0 \): mean rotation measure
- \( \sigma_{RM} \): width of \( RM \) distribution
- \( \sigma_I \): width of \( I \) distribution
- \( \sigma_{Q,U} \): width of \( Q, U \) distribution

### Model parameters that can be optimized

<table>
<thead>
<tr>
<th>Parameter</th>
<th>Set by dependence of</th>
</tr>
</thead>
<tbody>
<tr>
<td>( B_{reg,\parallel} )</td>
<td>( R \bar{M}<em>0 ) (( B</em>{reg,\parallel} ))</td>
</tr>
<tr>
<td>( B_{ran} )</td>
<td>(constant) strength of random ( B )-field</td>
</tr>
<tr>
<td>( B_{reg,\perp} )</td>
<td>perpendicular regular ( B )-field</td>
</tr>
<tr>
<td>( P_b )</td>
<td>polarized intensity of background</td>
</tr>
<tr>
<td>( \eta_b )</td>
<td>factor for depolarization of background</td>
</tr>
<tr>
<td>( \eta_b )</td>
<td>( \sigma_{Q,U} ) (( B_{ran}, B_{reg,\perp}, P_b ))</td>
</tr>
</tbody>
</table>

### Additional constraints:

- Background depolarization factor \( 0 \leq \eta_b \leq 1 \)
- Number of cells \( N = L/d \), while \( N f \) cells determine the shape of \( RM \) distribution
- \( B_{reg,\perp}/P_b \) determines shape of \( Q, U \) distribution

| \( B_{ran,\perp}/P_b \) | \( \bar{M}_0 \) (\( B_{ran,\perp} \))                                           |

### Table 3.4: Parameters in the depth depolarization model.

The first set of parameters is determined by literature or can be arbitrarily chosen. The second set of parameters is free to choose and is varied in the models. The third set of parameters is set by our observations. The last set are those parameters of the ISM that can be estimated from the models, followed by the input parameters from the categories above (mostly observational constraints) that determine the value of these parameters. In parentheses the model parameters that they depend on. These dependences are such that the order of parameter determination given in the table is compulsory. In this way it is possible to derive definite unambiguous values or small ranges of the five parameters from the five observables.

#### 3.9.2 The various types of model parameters

Four types of parameters are used in the model, which are shown in Table 3.4. We discuss these categories separately.
Input parameters with fixed values

These are physical parameters which are known or for which exist reasonably good estimates in the literature. From dispersion measures (DM) of pulsars in globular clusters at high Galactic latitude and Hα emission measures (EM), Reynolds (1991) derives \( n_e \approx 0.08 \text{ cm}^{-3} \) in the cells, with a filling factor \( f = 40\% \) if the warm ionized ISM layer has a constant electron density, and 20\% if the electron density distribution is exponential. The Beuermann et al. (1985) model for Galactic synchrotron radiation predicts a scale height of the thin disk of 180 pc. We run the model with a fixed \( f = 20\% \), \( h = 180 \text{ pc} \) and \( n_e = 0.08 \text{ cm}^{-3} \) and discuss afterwards how the results would change if these parameters were different.

The intrinsic polarization angle of the background only defines the average angle in the final map of polarization angle. It changes the \( Q \) and \( U \) maps locally, but has no influence on the distributions of \( Q \) and \( U \). Therefore the value of \( \phi_0 \) is arbitrary and chosen to be 0°. The angles \( \alpha \) and \( \phi_i \) define the orientation of the random component of the magnetic field, with respect to the line of sight and some fixed direction in the plane of the sky, respectively. Both are drawn from uniform distributions, for each individual cell.

The total intensity \( I_0 \) is taken from the 408 MHz all-sky survey by Haslam et al. (1982). The 2.7 K contribution from the CMBR is subtracted from the values in the survey before these are converted to a frequency of 350 MHz using a spectral index of \( \approx 2.7 \). Approximately 25\% of the total background temperature is due to point sources (from source counts, Bridle et al. 1972), so only the remaining 75\% is included in the model. This yields values of 34 K and 47 K for \( I_0 \) in Auriga and Horologium, respectively.

Free input parameters

No external constraints are imposed on cell size \( d \) and proportionality factor \( C \). Estimates of cell size are found in the literature (Rand and Kulkarni 1989, Ohno and Shibata 1993) but they range from 10 pc to several hundreds of pc, and mostly probe scales that exceed the size of our fields. A turbulent cell larger than a few degrees (\( \approx 20 \text{ pc} \) at a distance of 500 pc) will produce a constant component in our model. We probe cell sizes from a parsec to several tens of parsecs, and find the cell size determined in a reasonably narrow range because of the observational constraints.

The proportionality constant \( C \) is not free to choose but physically determined by the density of cosmic ray electrons. However, as the local cosmic ray density is not known, we treat \( C \) as a free parameter. After modeling, we compare the synchrotron emissivity from our model with estimates from the literature.

If strict equipartition between cosmic rays and magnetic field applies, then \( C \) is not constant but varies with \( B_z^2 \), so that the synchrotron emission \( I \propto B_z^2 \). Although equipartition is believed to hold on global Galactic scales, it is highly uncertain if equipartition is valid at parsec scales as well. Therefore, the exponent \( \alpha \) of the relation \( I \propto B_z^2 \) could be between 2 and 4, but its value does not influence the depolarization by a large factor, in the case of a source with a small-scale random magnetic field (Burn 1966). We assume \( \alpha = 2 \) in the model, and discuss afterwards the change in parameter
<table>
<thead>
<tr>
<th>Parameter</th>
<th>Auriga region</th>
<th>Horologium region</th>
</tr>
</thead>
<tbody>
<tr>
<td>$R M_0$</td>
<td>-3.4 rad m$^{-2}$</td>
<td>-1.4 rad m$^{-2}$</td>
</tr>
<tr>
<td>$\sigma_{R M}$</td>
<td>1.8 rad m$^{-2}$</td>
<td>4.3 rad m$^{-2}$</td>
</tr>
<tr>
<td>$\sigma_I$</td>
<td>$\leq 1.7$ K</td>
<td>$\leq 2.5$ K</td>
</tr>
<tr>
<td>$\sigma_{Q, U}$</td>
<td>3 K</td>
<td>4.3 K</td>
</tr>
</tbody>
</table>

**Constraints from the observations:**

Distributions of $R M$, $I$, $Q$, and $U$ are Gaussian.

Table 3.5: Values of observationally determined parameters from polarization maps of the Auriga and Horologium regions and other observational constraints for the models.

values if $\alpha > 2$.

**Constraints determined from the observations**

As discussed in Section 3.3, our observations yield distributions of $I$, $Q$, $U$, and $R M$. We summarize these results in Table 3.5, in the form of the mean value of observed $R M$, $R M_0$, and the dispersions in $R M$, $I$, $Q$, and $U$. In Auriga, the dispersion in $R M$ was derived after subtraction of a gradient in $R M$ (see Section 3.3). The mean values of $I$, i.e., $I_0$, were derived from external data, as discussed above. Each of the regions already provides useful constraints for these model parameters that can be freely adjusted and optimized (see below) but the combined set of constraints in Table 3.5 is quite powerful by virtue of the different path lengths through the medium.

**Model parameters that can be adjusted and optimized**

The model contains five parameters that are not derived from external data or from the observations. These are: the parallel and perpendicular components of the large-scale magnetic field, $B_{R g, \parallel}$ and $B_{R g, \perp}$ respectively, the strength of the random component of the magnetic field $B_{r a n}$, the uniform intensity of the polarized emission from the thick disk $P_b$, and the overall depolarization factor $\eta_b$ which connects $P_b$ and $I_b$ (see Section 3.9.1).

In adjusting and optimizing these five parameters, the constraints from the observations must be used in a specific order, as follows. The observed average rotation measure $R M_0$ depends only on $B_{R g, \parallel}$, so $B_{R g, \parallel}$ is set to fit the observed $R M_0$. The observed dispersion in $R M$, $\sigma_{R M}$, depends on $B_{R g, \parallel}$ and $B_{r a n}$, but because we derived $B_{R g, \parallel}$ from $R M_0$, only $B_{r a n}$ can be tuned to obtain the observed width $\sigma_{R M}$. Subsequently, with $B_{R g, \parallel}$ and $B_{r a n}$ determined, $B_{R g, \perp}$ and $P_b$ form a coupled parameter set. An increase in either of them causes an increase in both $\sigma_{Q, U}$ and $I_0$, but $\sigma_I$ depends only on $B_{R g, \perp}$ and not on $P_b$. Therefore either a maximum value of $B_{R g, \perp}$ follows from the observed upper limit to $\sigma_I$, with $P_b$ chosen to produce the correct value of $\sigma_{Q, U}$, or a range in $B_{R g, \perp}$ and $P_b$ is determined to match the observed value of $\sigma_{Q, U}$, while $\sigma_I$ remains consistent with the observational upper limit. The maximum value of $P_b$ is determined from the highest value of $P_b/B_{R g, \perp}$ for which the modeled $Q$ and $U$ distributions are still Gaussian. Finally, the parameter with which to adjust $I_0$ to
Figure 3.17: Dependences of depth depolarization parameters on observables in the Auriga region, for model A. Left plot: \( R\theta_0 \) increases with \( B_{\text{reg},||} \). Center plot: for fixed \( B_{\text{reg},||} \), \( \sigma_{RM} \) increases with \( B_{\text{ran}} \), with only a weak dependence on the value of \( B_{\text{reg},||} \). Right plot: \( \sigma_{Q,U} \) depends on \( B_{\text{ran}} \) and \( B_{\text{reg},\perp} \).

its desired value is \( \eta_0 \), all others already being determined in earlier steps. So due to the different dependences of observables on the parameters, definite values for some of these parameters can be found, instead of only an allowed parameter range. In some model runs, the desired value of \( \eta_0 \) is attained before reaching the maximum \( \sigma_f \). In that case, \( \sigma_f \) is left smaller than the maximum and \( B_{\text{reg},\perp} \) and \( P_b \) are constrained by \( \eta_0 \).

An example of the correlations that we used is given in Fig. 3.17. The leftmost plot shows the dependence of \( R\theta_0 \) on \( B_{\text{reg},||} \). As expected, a large regular magnetic field component causes a large non-zero mean rotation measure. (We have chosen a negative value of the magnetic field because the observed \( R\theta_0 \) is negative, see Table 3.5.) Then, for three fixed values of \( B_{\text{reg},||} \), the center plot gives the change of \( \sigma_{RM} \) with \( B_{\text{ran}} \), which increases roughly linearly with \( B_{\text{ran}} \) and shows hardly any dependence of \( \sigma_{RM} \) on \( B_{\text{reg},||} \). Having set \( B_{\text{ran}} \) to obtain the observed value of \( \sigma_{RM} \), the right plot shows how the observable \( \sigma_{Q,U} \) depends on \( B_{\text{reg},\perp} \). The width of the \( Q \) and \( U \) distribution depends slightly on the chosen values of \( B_{\text{ran}} \).

3.10 Comparison of model predictions with observations

For models A and B, as defined in Section 3.9.1, the propagation of polarized radiation through the medium is computed for a range of values of the adjustable parameters, viz. cell size \( d \), and \( C \), the proportionality factor between synchrotron emissivity and \( B_0^2 \), as well as the five parameters that can be adjusted and optimized by applying the observational constraints.

The allowed cell size is well-constrained by the observations: if the cell size is large, the number of cells is small for a given path length and filling factor. As the number of cells with Faraday-rotating, thermal medium can differ per line of sight, the \( RM \) distribution will not be Gaussian anymore if the regular \( RM \) component dominates,
i.e. $B_{reg||}$ is large. On the other hand, if the cell size is small, the $RM$ per cell decreases. But to obtain a large enough $\sigma_{RM}$, the $RM$ per cell has to be rather high, so the parameter $B_{ran}$ has to be increased to produce the observed value of $\sigma_{RM}$. However, an increase of $B_{ran}$ increases $\sigma_I$, which then puts an upper limit on $B_{reg,\perp}$. To produce the observed dispersion in $Q$ and $U$, we then need a large value for the background polarized intensity $P_b$. If the cell size is taken too small, $P_b$ becomes so large compared to the polarized emission in the cells in layer I, that the distributions of $Q$ and $U$ become distinctly non-Gaussian, in disagreement with the observations. Allowed values of $d$ range from approximately 1 to 60 pc, with an optimum value of about 15 pc, in good agreement with estimates by Ohno and Shibata (1993). However, a cell size of 15 pc located at the far end of the thin disk in the direction of the Auriga region subtends an angle of more than a degree on the sky. As we observe structure on degree to arcminute scales, smaller cells must be present as well. Most likely, cell sizes in a wide range are present on these scales (Minter and Spangler 1996, Clegg et al. 1992). However, note that the computed cell size gives the scale over which $RM$ varies, while the scale size over which the polarization angle changes by $90^\circ$ is also essential for beam depolarization, and can be different from the cell size computed above.

The constant $C$, the synchrotron emission per cell for a given value of $B_0^2$, is not directly constrained. However, its only influence on the data is changing $B_{reg,\perp}$ (and therefore $B_{ran}/B_{reg}$), which varies from $0.3 \, \mu G \geq B_{reg,\perp} \leq 27 \, \mu G$ for $20 > C > 0.01$. We use the intermediate value $C = 1$ in the models, keeping in mind that values for $B_{reg,\perp}$ are allowed to vary.

We now summarize the result of the comparison between the models and the observations for the allowed ranges of $B_{ran}$, $P_b$, and $\eta_b$. The cell sizes probed in the modeling were 1, 2, 5, 10, 20, 30, 40 and 60 pc, although for model B, only cell sizes above 5 pc are allowed, and smaller cell sizes are allowed for the Auriga region than for Horologium.

Fig. 3.18 shows the allowed ranges of parameters for models A and B for the two regions, for $C = 1$. The upper plots show values of $B_{reg||}$ obtained in model A (left) and model B (right), where the solid line denotes the Auriga region and the dotted line the Horologium region. The parallel regular field $B_{reg||}$ is about $-0.42 \pm 0.02 \, \mu G$ for Auriga and $-0.085 \pm 0.005 \, \mu G$ for Horologium in model A, and $-0.35 \pm 0.01 \, \mu G$ and $-0.065 \pm 0.005 \, \mu G$ respectively in model B, where the errors are estimated from the spread in observed values. $B_{reg||}$ hardly depends on $C$ or $d$. The plots below these show $B_{ran}$, where the best value is about 1 $\mu G$ for large ($\geq 5$ pc) cell sizes for both the Auriga and the Horologium region, and increases to $\sim 4 \, \mu G$ for cell sizes of a parsec. $B_{ran}$ increases with decreasing cell size because, if the path length through a single cell is smaller, a higher $B_{ran}$ is needed to create the observed width of the $RM$ distribution.

For the remaining parameters $B_{reg,\perp}$, $P_b$, and $\eta_b$, only parameter ranges could be determined instead of definite values, given in Fig. 3.18 by dark grey for Auriga and light grey for Horologium. The edges of the shaded areas are not sharp but have a considerable uncertainty. The allowed parameter range should be read more as an indication of possible parameter values than as ranges with sharp boundaries. Moreover, the discrete edges at a certain cell size are not physical, but only indicate that the next probed cell size could not produce parameters in agreement with the observations.
Figure 3.18: Allowed ranges of parameters for model A (left) and B (right) for different values of cell sizes and $C = 1$. The magnetic field values are given in $\mu G$, the polarized brightness temperature of the background $P_b$ in K and $I_{0,\text{thin}}$ is the percentage of the total emissivity that originates in the thin disk. The lines in the upper plots show values found for $B_{\text{reg},1}$ and $B_{\text{reg},2}$ for the Auriga region (solid line) and the Horologium region (dotted line). All plots below those show allowed ranges in parameters, given in dark grey for the Auriga region and light grey for Horologium.

A basic first conclusion is that the values obtained in the two regions roughly agree, even though the Auriga and Horologium regions have different input parameters and a different line of sight through the medium. The regular magnetic field components approximately agree in the Auriga and Horologium regions ($B_{\text{reg}} \approx 3 \mu G$). The perpendicular and the parallel component of the regular magnetic field are almost constant over cell size, approximately $2.8 \pm 0.5 \mu G$ in Auriga and $3.3 \pm 0.5 \mu G$ in Horologium for model A, and $2.8 \pm 0.5 \mu G$ in Auriga and $3.1 \pm 0.5 \mu G$ in Horologium for model B. The intensity of the polarized background $P_b$ varies between 0.1 and 3 K, with a best
estimate of about 1.5 ± 1.0 K. This is sufficiently small that the distributions of Stokes $Q$ and $U$ can be quite close to Gaussian. The factor $\eta_b = P_b/(0.7 I_b)$ ranges from almost zero to 0.6. Our best estimate is about 0.15 ± 0.1. At 350 MHz, depolarization of the constant background to 0.15 times the original polarization is caused by $RM \approx 3 - 5 \text{ rad m}^{-2}$ (Burn 1966), which indicates a value of $n_eB_\parallel \approx 0.1 - 0.2 \mu \text{G cm}^{-3}$ for a height of the Reynolds layer of 1 kpc. Thus, for $n_e \approx 0.05 \text{ cm}^{-3}$ in the halo, the halo magnetic field could persist without much attenuation throughout the Reynolds layer, as was suggested earlier by Han et al. (1999).

Our estimate of $B_{\text{ran}}/B_{\text{reg}} \leq 1$ is lower than most of the estimates from the literature discussed in Section 3.8. This may be due to several factors. First, our regions of observation are quite large (7° - 9° on a side) but large-scale random magnetic field could still have been interpreted as regular field in our analysis. Secondly, it could be the result of selection, as our observational fields were chosen for their high polarized intensity, which in our model automatically implies a modest random magnetic field. Finally, they are in the second Galactic quadrant, so we probe mostly the inter-arm region between the Local and the Perseus arms, where $B_{\text{ran}}/B_{\text{reg}}$ is smaller than in the average ISM including spiral arms (Han and Qiao 1994, Indrani and Deshpande 1998, Beck 2001).

The bottom plots in Fig. 3.18 give $I_{0,\text{thin}}$, the total synchrotron emission of the thin disk. The emission in the thin disk is also estimated by Beuermann et al. (1985) in their standard decomposition of $I_0$ into thin and thick disk contributions. According to their model, only about 20 to at most 35% of $I_0$ is generated in the thin disk and the nearest 180 pc of the thick disk. Furthermore, Caswell (1976) estimated the synchrotron emissivity from a survey with the Penticton 10 MHz array as 240 K pc$^{-1}$ at 10 MHz. Rescaled to 350 MHz, this gives a total emission from the thin disk of 10.6 K in the Auriga region and 21 K in the Horologium region. Roger et al. (1999) estimate from the 22 MHz survey performed with the DRAO 22 MHz radio telescope an emissivity of about 55 K pc$^{-1}$ for two HII regions in the outer Galaxy, out of the Galactic plane. Their results give estimates of the emissivity in the thin disk which are approximately twice as high as the estimates from the Caswell survey. Due to the large uncertainty in the emissivity, $I_{0,\text{thin}}$ does not put a strong constraint on the model parameters.

Having determined parameter ranges, we vary the set of parameters defined in Section 3.9.2 as parameters that are kept fixed at externally determined values, viz. electron density $n_e$, filling factor $f$ and height of the layer $h$. We choose a particular set of values for all earlier determined parameters, and vary filling factor $f$, thermal electron density $n_e$ or height $h$ to see their influence on the model parameters. A filling factor $f \leq 5 - 10\%$ is not allowed in either model: large cell sizes give a non-Gaussian $RM$ distribution, and small cell sizes yield too high a background polarization to keep $Q$ and $U$ Gaussian. No upper limit can be given for the filling factor, although $B_{\text{ran}}/B_{\text{reg}}$ decreases with a factor two for $f = 1$. For varying thermal electron density, a low $n_e \leq 0.03 \text{ cm}^{-3}$ dictates such a high $B_{\text{ran}}$ that the ratio $B_{\text{reg},\perp}/P_b$ becomes so low that $Q$ and $U$ become distinctly non-Gaussian. High electron densities are allowed in the models but the random magnetic field drops to very low values ($B_{\text{ran}} \leq 0.15 \mu \text{G}$ for $n_e \geq 0.1 \text{ cm}^{-3}$). A lower limit to the height of the thin disk is about 100 pc, again no upper limit can be set.
We checked the influence of the assumption of equipartition between magnetic field and cosmic rays on these scales. If $I \propto B^4$, instead of $I \propto B^2$, the upper limit to structure in $I$ becomes much more stringent. Therefore, model A will no longer produce any solutions that agree with the observables. In model B, where the random magnetic field is restricted to the cells with thermal electrons, solutions are found with cell sizes 10 and 20 pc, and $B_{\text{reg.1}}$ much lower, about 0.5 $\mu$G. Other parameters are comparable to the case where $C$ is constant.

Finally, it should be mentioned that the mean of the distributions of $Q$ and $U$, referred to as offsets in Section 3.4, is lower than 1.2 K in all models for all parameters, which is consistent with the discussion in Section 3.4.

3.10.1 Estimate of the polarization horizon

In a medium where depth depolarization occurs, polarized radiation originating at large distances from the observer will become depolarized more easily than emission originating nearby. So most of the observed polarized radiation is created in the nearest layer, and the distance out to which polarization is observed has been referred to as the “polarization horizon” (Landecker et al. 2001).

In our model, we build up a line of sight by adding cells one by one, starting at the observer. The radiation from each added cell is Faraday-rotated by all warm ionized material in front of it. The observed degree of polarization after addition of each cell is given as a function of the line of sight built up until that particular cell in Fig. 3.19, for the Auriga (solid line) and Horologium (dashed line) regions, for model A with a cell size of 5 pc. Still a fairly large fraction (20%) of the polarized emission from distant layers can be observed, so depth depolarization alone cannot produce a true horizon, but attenuates the polarization from distant layers gradually.

The second effect which creates a polarization horizon is increasing beam depolarization because structure at larger distances subtends a smaller angle on the sky. If the smallest scales in the observed regions are about a parsec, the angle of these scales on the sky becomes smaller than the beam at a distance of about 700 pc. Spoelstra (1984) derived the polarization horizon from comparison of radio continuum data at five frequencies from 408 MHz to 1411 MHz (Brouw and Spoelstra 1976) with starlight polarization. He found a distance to the origin of the polarized radio emission of $625 \pm 125$ pc in the direction of our fields, in agreement with our observations.
3.11 Conclusions

Small-scale structure in the linearly polarized component of the diffuse Galactic synchrotron emission is present in almost every direction. As a rule, this structure is not correlated with total emission, and therefore cannot be due solely to small-scale structure in emission. Instead, the polarization angle is Faraday-rotated by magneto-ionic material along the path length of the radiation. However, the structure in $P$ cannot be produced by Faraday rotation alone (which only rotates the polarization angle), but there are several other processes responsible for this. First, instrument-related effects produce structure in $P$, such as large-scale components in the radiation that are undetectable with an interferometer, depolarization due to variation in angle within the telescope beam, or over the frequency band width. Furthermore, physical depolarization processes in the ISM can cause depolarization if Faraday rotation and synchrotron emission occur in the same medium.

In this paper, we have discussed these processes and gauged their relative importance, using two sets of observations made with the Westerbork Synthesis Radio Telescope (WSRT). From the constraints imposed by the observations, we have derived estimates of several parameters of the warm ISM.

Undetectable large-scale components in Stokes $Q$ and/or $U$ measurements, so-called offsets, can create structure in polarized intensity, and prohibit the correct determination of rotation measures. However, we showed that in our fields, the observed range in rotation measure is so large that offsets cannot play a significant rôle.

Secondly, narrow, one-beam-wide canals of depolarization can be caused by beam depolarization or differential Faraday rotation. Although from our observations neither of the explanations can be excluded, beam depolarization is slightly favored. Most likely, both effects are present in the medium.

Beam depolarization is present in the observations (most notably in depolarization channels). It leaves a characteristic pattern in low-intensity regions as well. However, beam depolarization cannot explain all observed structure in the maps. Bandwidth depolarization is negligible in our measurements as the observed range in $RM$ is far too small to produce a significant angle change over the 5 MHz wide frequency band.

Depth depolarization, the depolarization process in a medium of synchrotron-emitting and Faraday-rotating material, is the dominant cause of structure in polarized intensity. We constructed a simple model of a layer of cells containing random and regular magnetic field $B_{\text{ran}}$ and $B_{\text{reg}}$, and a thermal electron density $n_e$ in a fraction $f$ of the cells (mimicking the filling factor $f$). This layer corresponds to the thin Galactic disk with small-scale structure in the magnetic field. The Galactic thick disk or halo is modeled by a constant background $P_h$, with a certain constant depolarization denoted by the factor $\eta_b$. We vary cell size, emissivity, magnetic fields and background to obtain a range of models that comply to the observational constraints, i.e. yield the correct width, mean and shape of the distributions of $Q$, $U$, $I$ and $RM$.

The results can be summarized as follows: the allowed cell size is constrained to be in the range of 1 to 60 pc, with a best estimate of 15 pc. The regular magnetic field component along the line of sight ($\sim -0.4 \, \mu G$ for Auriga, and $\sim -0.08 \, \mu G$ for Horologium) is much smaller than the regular magnetic field component perpendicular to the line of sight ($\sim 2.8 \, \mu G$ for Auriga, and $\sim 3.2 \, \mu G$ for Horologium), indicating
that the regular magnetic field is directed almost perpendicular to the line of sight in these directions. The random magnetic field component is about 1 to 3 $\mu$G in the two regions. In most models, the regular component of the magnetic field was higher than the random component, with an average ratio of $B_{\text{ran}}/B_{\text{reg}} = 0.7 \pm 0.5$ which increases for smaller cell sizes. Estimates from the literature tend to give larger ratios (0.5 - 4). This could be explained by the relatively small fields of view we use ($\sim 7^\circ$ in size), so that large-scale random components of the magnetic field are misinterpreted as regular components. Furthermore, the fields of observation were selected for their high polarization, indicating a higher regular magnetic field component than average, and are situated in an inter-arm region, where uniform magnetic fields tend to be higher than average. The constant polarized background intensity from the thick disk is about $1.5 \pm 1.0$ K. The attenuation of polarized emission at larger distances due to depth depolarization is not very strong (down to 20% at 1200 pc), so a 'polarization horizon' is not well-defined.

Acknowledgements

We wish to thank R. Beck and E. Berkhuijsen for their critical comments and helpful discussions. The Westerbork Synthesis Radio Telescope is operated by the Netherlands Foundation for Research in Astronomy (ASTRON) with financial support from the Netherlands Organization for Scientific Research (NWO). MH acknowledges support from NWO grant 614-21-006.

References

Beck R., 2001, SSRv 99, 243
Bridle A. H., Davis M. M., Fomalont E. B., Lequeux J., 1972, NPhS 235, 123
Gardner F. F., & Whiteoak J. B., 1966, ARAA 4, 245
Chapter 3

Indrani C., & Deshpande A. A., 1998, NewA 4, 331
Landecker T. L., Uyanik B., & Kothes R., 2001, AAS 199, #58.07
Seeger C. L., & Westerhout G., 1961, AJ 66, 294
Multi-frequency polarimetry of the Galactic radio background around 350 MHz: I. A region in Auriga around $l = 161^\circ$, $b = 16^\circ$

M. Haverkorn, P. Katgert and A. G. de Bruyn, submitted to A&A

Abstract

With the Westerbork Synthesis Radio Telescope (WSRT), multi-frequency polarimetric images were taken of the diffuse radio synchrotron background in a $\sim 5\degree \times 7\degree$ region centered on $(l, b) = (161^\circ, 16^\circ)$ in the constellation of Auriga. Observations were done simultaneously in 5 frequency bands, from 341 MHz to 375 MHz, and have a resolution of $\sim 5.0^\prime \times 5.0^\prime$ cosec $\delta$. Polarized intensity $P$ and polarization angle $\phi$ show ubiquitous structure on arcminute and degree scales, with polarized brightness temperatures up to about 13 K. On the other hand, no structure at all is observed in total intensity $I$ to an r.m.s. limit of 1.3 K. This indicates that the structure in the polarized radiation must be due to Faraday rotation and depolarization mostly in the warm component of the nearby Galactic ISM. Therefore, study of the polarized emission can yield information on structure in the warm ISM. Different depolarization processes create structure in polarized intensity $P$. Beam depolarization creates depolarization “canals” of one beam wide, while depth depolarization is thought to be responsible for creating most of the structure on scales larger than a beam width. The observations are interpreted in terms of a simple model of the warm ISM with a polarized background. In that model the warm ISM consists of cells of a single size containing relativistic and thermal gas and regular and random magnetic field, in which synchrotron emission and Faraday rotation is produced. The observations are best fitted with a cell size of 10 to 20 pc and a ratio $B_{\text{reg}}/B_{\text{reg}} \approx 0.7 \pm 0.5$. The polarization horizon, beyond which most diffuse polarized emission is depolarized, is estimated to be a distance of about 600 pc. Rotation measures ($RM$) can be reliably determined, and are in the range $-17 \lesssim RM \lesssim 10$ rad m$^{-2}$ with an non-zero average $RM_0 \approx -3.4$ rad m$^{-2}$. The distribution of $RM$s on the sky shows both abrupt changes on the scales of the beam and a gradient in the
direction of positive Galactic longitude of \( \sim 1 \text{ rad m}^{-2} \) per degree. The gradient and average \( R M \) are consistent with a regular magnetic field of \( \sim 1 \mu G \) which has a pitch angle of \( p = -14^\circ \). There are 13 extragalactic sources in the field for which \( R M s \) could be derived, and those have \( |R M| \leq 13 \text{ rad m}^{-2} \), with an estimated intrinsic source contribution of \( \sim 3.6 \text{ rad m}^{-2} \). The \( R M s \) of the extragalactic sources show a gradient that is about 3 times larger than the gradient in the \( R M s \) of the diffuse emission and that is approximately in Galactic latitude. This difference is ascribed to a vastly different effective length of the line of sight. The \( R M s \) of the extragalactic sources also show a sign reversal which implies a reversal of the magnetic field across the region on scales larger than about ten degrees.

4.1 Introduction

The warm ionized gas and magnetic field in the Galactic disk Faraday-rotate and depolarize the linearly polarized component of the Galactic synchrotron emission. Structure of the warm ionized Interstellar Medium (ISM) is imprinted in the polarization angle and polarized intensity by Faraday rotation and depolarization. Therefore, the structure in the polarized diffuse radio background provides unique information about the structure in the magneto-ionized component of the ISM. Using interferometers we can probe arcminute scales, which correspond to linear scales from a fraction of a parsec to several tens of parsecs. Multi-frequency radio polarization observations of the Faraday-rotated synchrotron emission enable determination of the rotation measure (\( R M \)) distribution along many contiguous lines of sight. Therefore, this is a valuable method to estimate the non-uniform component of the Galactic magnetic field, weighted with electron density.

Although the magnetic field is only one of the factors that shape the very complex multi-component ISM, it is generally thought to play a major rôle in the energy balance of the ISM and in setting up and maintaining the turbulent character of the medium. The non-uniform component of the magnetic field may influence heating of the ISM (Minter and Balser 1997), and provide global support of molecular clouds (e.g. Vázquez-Semadeni et al. 2000, Heitsch et al. 2001). It also plays an important rôle in star formation processes (e.g. Shu 1985, Beck et al. 1996, Ferrière 2001). Furthermore, observations of the detailed structure of the magnetic field can provide constraints for Galactic dynamo models (Han et al. 1997).

The strength and structure of the magnetic field in the warm ionized medium in the Galaxy can also be obtained from Faraday rotation measurements of pulsars, or polarized extragalactic sources.

Among these, radio observations of pulsars take a special place because for them one can measure the \( R M \) and the dispersion measure (\( DM \)) along the same line of sight. The ratio of the two immediately yields the averaged component of the magnetic field along a particular line of sight (see e.g. Lyne and Smith 1989, Han et al. 1999). These studies, and analyses of the distribution of polarized extragalactic sources (e.g. Simard-Normandin and Kronberg 1980, Sofue and Fujimoto 1983), suggest that the regular Galactic magnetic field is directed along the spiral arms, with a pitch angle of \( p \approx 5 - 15^\circ \). There is evidence for reversals of the uniform magnetic field inside the solar
Multi-frequency polarimetry around 350 MHz I

circle (e.g. Simard-Normandin and Kronberg 1979, Rand and Lyne 1994), while there is still debate on reversals outside the solar circle (e.g. Vallée 1983, Brown and Taylor 2001, Han et al. 1999). Rand and Kulkarni (1989) used pulsar $RMs$ to derive a value for the strength of the regular component of the magnetic field of $B_{\rm reg} = 1.3 \pm 0.2 \, \mu G$. This result assumes a circularly symmetric large-scale Galactic magnetic field. The residuals with respect to the best-fit large-scale model were interpreted as due to a random magnetic field of $5 \, \mu G$. This number was derived by assuming a scale length of the random field, modeled as a cell size, of 55 pc. Ohno and Shibata (1993) confirmed this result for the random magnetic field component for all cell sizes in the range of 10 - 100 pc. They used 182 pulsars in pairs and therefore did not have to make any assumptions about a large-scale Galactic magnetic field.

Using extragalactic sources, most of which are double-lobed with lobe separations between $30''$ and $200''$, Clegg et al. (1992) found substantial fluctuations in $RM$ on linear scales of $\sim 0.1$ - $10$ pc, which they could explain by electron density fluctuations alone. However, Minter and Spangler (1996) observed $RMs$ fluctuations in extragalactic source components that cannot be explained by only electron density fluctuations; they need an additional turbulent magnetic field of $\sim 1 \, \mu G$ to fit the observations. Observations of polarization of starlight (Jones et al. 1992) give much larger estimates of the cell size, of up to a kpc.

Using pulsars and extragalactic radio sources for the determination of the $RM$ of the Galactic ISM is clearly not ideal, because they only provide information in particular directions which sample the ISM very sparsely. On the contrary, the diffuse Galactic radio background provides essentially complete filling over large solid angles. Therefore, rotation measure maps of the diffuse emission can be produced that give information on the electron-density-weighted magnetic field over a large range of scales. The distribution of polarized intensity, when interpreted as mostly being due to depolarization, can yield estimates of several properties of the warm ISM such as correlation length, ratio of random over regular magnetic field and the distance out to which diffuse polarization can be observed. Early $RMs$ maps of the diffuse synchrotron emission in the Galaxy were constructed by Bingham and Shakeshaft (1967) and Brown and Spoelstra (1976) who confirmed a Galactic magnetic field in the Galactic plane. Junkes et al. (1987) presented a polarization survey of the Galactic plane at $4.9^\circ < l < 76^\circ$ and $|b| < 1.5^\circ$ showing small-scale structure in diffuse polarization. The existence of polarization filaments at intermediate latitudes, without correlated structure in total intensity $I$, was discovered by Wieringa et al. (1993) at 325 MHz. Polarization surveys have been performed at frequencies from 1.4 GHz to 2.695 GHz (Duncan et al. 1997, Duncan et al. 1999, Uyaniker et al. 1999, Landecker et al. 2001, Gaensler et al. 2001), mostly in the Galactic plane.

In this chapter, we discuss the results of multi-frequency observations at low frequencies around 350 MHz of a field in the constellation Auriga, in the second Galactic quadrant ($l = 161^\circ$, $b = 16^\circ$). Due to the low frequencies, we probe low rotation measures that are predominant at intermediate and high latitudes. This gives the opportunity to study the high-latitude $RM$ without concrete objects such as HII regions or supernova remnants in the line of sight, and estimate structure in the ISM above the thin stellar disk.

Section 4.2 contains details of the multi-frequency polarization observations. In
Section 4.3, we analyze the observations, in particular the small-scale structure in polarized intensity and polarization angle. Faraday rotation is discussed in Section 4.4, where we also present the map of rotation measure. In Section 4.5, depolarization mechanisms are described that cause the structure in $P$, and the constraints that the observations provide for the parameters that describe the warm ISM. In Section 4.6, the polarization properties of 13 polarized extragalactic point sources found in the Auriga region are discussed. In Section 4.7, we discuss the information that our data provides on the strength and structure of the Galactic magnetic field. Finally, our conclusions are stated in Section 4.8.

4.2 The observations

We used the Westerbork Synthesis Radio Telescope (WSRT) for multi-frequency polarimetric observations of the Galactic radio background in a field in the constellation of Auriga centered on $(\alpha, \delta)$ (B1950) = (6$^h$20$^m$, 52°30$^m$) ($l = 161^\circ$, $b = 16^\circ$). This field was observed in 8 frequency bands between 325 and 300 MHz simultaneously, where the width of each frequency band was 5 MHz. Due to radio interference and hardware problems we could only use the data in 5 of the 8 bands, viz. those centered on 341, 349, 355, 360, and 375 MHz.

The region in Auriga was observed in six 12hr periods, which resulted in a baseline increment of 12m. The shortest baseline obtained is 36m, and the longest is 2700m, which gives a maximum resolution of ~ 1$''$. For the study of the diffuse polarized emission we tapered the $(u,v)$-data to increase the signal-to-noise ratio, so as to obtain a resolution of $5.0' \times 5.0' \cos \delta = 5.0' \times 6.3'$ in all 5 frequency bands. As an interferometer has a finite shortest baseline, it is insensitive to large-scale structure. The shortest spacing of 36m in the WSRT constitutes effectively a high-pass filter for all scales above approximately a degree.

The Auriga region was selected from diffuse polarization maps that were produced as a by-product of the Westerbork Northern Sky Survey (WENSS, Rengelink et al. 1997), which is a single-frequency radio survey at 325 MHz. The WENSS diffuse polarization maps contain several regions of high polarization which show conspicuous small-scale structure in polarized intensity and polarization angle. We reobserved two of those regions at multiple frequencies and with higher sensitivity to obtain rotation measure information. The first region in Auriga is described in this chapter, and the other one in Chapter 5. In addition, we have used the single-frequency WENSS polarization data to construct a large polarization map of about $30' \times 35'$ to study the large-scale features of the polarized radio background (see Chapter 7).

The data reduction process is described in detail in Chapter 2, and we only give a brief summary here. The observations were reduced using the NEWSTAR data reduction package. Polarized and unpolarized standard calibrator sources were used, where the absolute flux scale at 325 MHz is based on a value of 26.93 Jy for 3C286 (Baars et al. 1977). From this value the flux scales of the other calibrator sources 3C48, 3C147, 3C345, and 3C303 were derived.

As the area to be mapped is larger than the primary beam of the WSRT, the mosaicking technique was used (Rengelink et al. 1997). In this observing mode, the
Auriga field

| Central position | $|\text{l}_b| = (161^\circ, 16^\circ)$ |
| Size | $\sim 7^\circ\times 9^\circ$ |
| Pointings | $5 \times 7$ |
| Frequencies | $341, 349, 355, 360, 375 \text{ MHz}$ |
| Resolution | $5.0' \times 5.0' \csc \delta = 5.0' \times 6.3'$ |
| Noise | $\sim 4 \text{ mJy/beam (0.5 K)}$ |
| Conversion Jy$^{-1}$ K$^{-1}$ | $1 \text{ mJy/beam} = 0.127 \text{ K (at 350 MHz)}$ |
| Spacings | 36m, 48m, 60m, 72m, 84m, 96m |
| Observing date | 96/02/19, 95/12/24, 95/12/31, 96/01/07, 96/01/30, 96/01/09 |
| Start time (UT) | 13:36, 17:28, 17:13, 16:37, 14:59, 16:33 |
| End time (UT) | 01:54, 05:41, 05:31, 04:55, 03:12, 04:46 |

Table 4.1: Observational details of the Auriga field

array cycles through a preselected set of pointing positions a number of times during the 12 hour observation period, which reduces the instrumental polarization in between pointing centers to below 1%, see Chapter 2. For the present region we used $5 \times 7$ pointing positions separated by $1.25^\circ$. We only discuss the central $\sim 7^\circ \times 9^\circ$ of the observed field of view, as the edges of the mosaic exhibit larger noise due to primary beam attenuation and increased instrumental polarization. Maps were constructed of Stokes parameters $I$, $Q$, $U$, and $V$. From Stokes $Q$ and $U$, polarized intensity $P$ and polarization angle $\phi$ were derived as in Eqs. (2.9) and (2.10).

To avoid spurious solar emission coming in through polarized side lobes, all observations were done at night, as shown in Table 4.1. In addition, the observing period was close to solar minimum and in winter, minimizing ionospheric Faraday rotation. The high polarized brightness that we observe (see Section 4.3.2) indicates that ionospheric Faraday rotation does not vary much during each of the 12hr observations, because if it did, hardly any polarized intensity would have resulted. The differences between the average ionospheric Faraday rotation in the six 12hr observations were determined from several point sources (in the $Q$- and $U$-maps at $\sim 1'$ resolution). The result is shown in Fig. 4.1 for 4 strong sources and 4 weaker sources. The strong sources have a signal-to-noise $\sigma = 2 - 3.5$ for an individual 12hr observation, which corresponds to an uncertainty in polarization angle $\sigma_{\phi} \approx 10^\circ$. The signal-to-noise of the weaker sources is $\sigma = 1.2 - 1.5 (\sigma_{\phi} \approx 22^\circ)$. From Fig. 4.1 we conclude that differential Faraday rotation between the six 12hr observations is negligible to within the uncertainties with which it can be measured from the point sources.

We can estimate the $RM$ contribution of the ionosphere from the earth magnetic field and the electron density in the ionosphere. The contribution of the ionospheric Faraday rotation is estimated as follows. The total electron content (TEC) in the ionosphere above Westerbork at night, in a solar minimum, and in winter is minimal: $\text{TEC} \approx 2.2 \times 10^{16}$ electrons cm$^{-2}$ (Campbell, private communication). The Auriga field at declination $\delta = 52.5^\circ$ is almost in the zenith at Westerbork at hour angle zero. If
we assume a vertical component of the earth magnetic field of 4.5 G towards us, and a path length through the ionosphere of 300 km, the \( RM \) caused by the ionosphere is \(-0.25 \text{ rad m}^{-2}\) at hour angle zero. At larger hour angles, this is even less. So we expect the rotation measure values given in this chapter not to be affected by ionospheric Faraday rotation by more than 0.5 rad m\(^{-2}\). No corrections were applied.

### 4.2.1 Missing large scale structure

An interferometer is increasingly insensitive to structure on larger angular scales due to missing small spacings. This means for the WSRT that scales above approximately a degree cannot be detected. The \( Q \) and \( U \) maps are constructed so that in each mosaic pointing, the map integral of \( Q \) and \( U \) are zero. This leads to missing large-scale components in \( Q \) and \( U \), and therefore erroneous determinations of \( P \), \( \phi \) and \( RM \). However, the distribution of \( RM \) is broad enough within one pointing. Therefore, the variation in polarization angle is so large that the average \( Q \) and \( U \) are close to zero, so that missing large-scale components are negligible (Section 4.5).

### 4.3 Total intensity and linear polarization maps

#### 4.3.1 Total intensity

In the left-hand panel of Fig. 4.2 we show the map of total intensity \( I \) at 349 MHz with a resolution of 5.6\('\times6.3\)'\'. Point sources were removed down to the confusion limit of 5 mJy/beam. Total intensity \( I \) has a Gaussian distribution around zero with width \( \sigma_T = 1.3 \text{ K} \) (while the noise, computed from \( V \) maps, is \( \sigma_N = 0.5 \text{ K} \)). \( I \) is distributed around zero due to the insensitivity of the interferometer to scales larger than approximately a degree. From the continuum single-dish survey at 408 MHz (Haslam et al. 1981, 1982), the \( T \)-background at 408 MHz in this region of the sky is estimated to be \( \sim 33 \text{ K} \) with a temperature uncertainty of \( \sim 10\% \) and including the 2.7 K contribution of the cosmic microwave background. Assuming a temperature spectral index of
Figure 4.2: Map of the total intensity $I$ (left) and polarized intensity $P$ (right) at 349 MHz, on the same brightness scale. The maps are smoothed to a resolution of $5.0' \times 6.3'$. White denotes high intensity, up to a maximum of $\sim 13$ K. Superimposed in the right map are lines of constant Galactic longitude, at Galactic latitudes of $13^\circ$, $16^\circ$ and $19^\circ$.

$-2.7$, the backgrounds at the lowest (341 MHz) and the highest (375 MHz) frequency of observation are $\sim 49$ K and $\sim 38$ K, respectively. Of these, approximately 25% is due to sources, as estimated from source counts (Bridle et al. 1972), and assuming the spectral index of the Galactic background and the extragalactic sources to be identical. We thus estimate the temperatures of the diffuse Galactic background at 341 and 375 MHz to be 37 and 29 K, respectively.

4.3.2 Polarized intensity

In the right-hand panel of Fig. 2 we show the map of polarized intensity at 349 MHz, with a resolution of $5.0' \times 6.3'$. The noise in this map is $\sim 0.5$ K, and the polarized brightness temperature goes up to $\sim 13$ K, with an average value of $\sim 2.3$ K.

The map shows cloudy structure of a degree to a few degrees in extent, with linear features that are sometimes parallel to the Galactic plane. In addition, a pattern of black narrow wiggly canals is visible (see e.g. the canal around $(\alpha, \delta) = (92.7^\circ, 49^\circ - 51^\circ)$). These canals are all one synthesized beam wide and have been shown to be due to beam depolarization. They separate regions of fairly constant polarization angle between which the difference in polarization angle is approximately $90^\circ$ ($\pm n 180^\circ$, $n = 1, 2, 3 \ldots$) (Haverkorn et al. 2000, Chapters 3 and 6). A change of $90^\circ$ (or $270^\circ$, $-90^\circ$)
Table 4.2: Correlation coefficients for correlations between $P$ and $I$, between $I$ in different frequency bands and between $P$ in different frequency bands, for tapered data and full resolution data.

\[
\begin{array}{|c|c|c|c|c|}
\hline
\text{freq band} k \quad \text{MHz} & \text{res } \tau & C(I_k, P_k) & C(P_{341}, P_k) & C(I_{341}, I_k) \\
\hline
341 & 5 & -0.008 & 1.00 & 1.00 \\
349 & 5 & -0.008 & 0.67 & 0.45 \\
355 & 5 & -0.011 & 0.65 & 0.34 \\
360 & 5 & 0.002 & 0.58 & 0.31 \\
375 & 5 & 0.015 & 0.53 & 0.24 \\
\hline
341 & 1 & 0.007 & 1.00 & 1.00 \\
349 & 1 & 0.004 & 0.60 & 0.08 \\
355 & 1 & 0.0004 & 0.60 & 0.21 \\
360 & 1 & 0.0002 & 0.53 & 0.15 \\
375 & 1 & 0.005 & 0.50 & 0.16 \\
\hline
\end{array}
\]

450° etc.) in polarization angle within one beam cancels the polarized intensity in the beam; therefore these canals appear black in Fig. 4.2. Hence, the canals are not physical structures, but the reflection of specific features in the distribution of polarization angle. Such canals have also been detected by others at higher frequencies (e.g. Uyannker et al. 1999, Gaensler et al. 2001). It must be appreciated that the canals represent an extreme case of beam depolarization. Beam depolarization is much more wide-spread if the polarization angle varies on the scale of the beam or on smaller scales, but in general its effect is (much) less than in the canals. Therefore, with the exception of the canals, beam depolarization does not leave easily visible traces in the polarized intensity distribution.

There is no obvious structure in Stokes $I$, and if there is any, it does not appear to be correlated with the structure in $P$. This is not typical for the Auriga field, but it is true in all fields observed so far with the WSRT at ~ 350 MHz (Katgert and de Bruyn 1999, and Chapters 5 and 7).

We have estimated the amount of correlation between total and polarized intensity by deriving the correlation coefficient $C$, where $C(f, g)$ of the observables $f(x, y)$ and $g(x, y)$ is defined as

\[
C(f, g) = \frac{\sum_{i,j}(f_{ij} - \overline{f})(g_{ij} - \overline{g})}{\sqrt{\sum_{i,j}(f_{ij} - \overline{f})^2 \sum_{i,j}(g_{ij} - \overline{g})^2}} \tag{4.1}
\]

where $\overline{f}$ is the average value of $f(x, y)$, $f_{ij} = f(x_i, y_j)$, summation is over $i = 1, N_x$ and $j = 1, N_y$, and $N_x$ and $N_y$ is the number of data points in the $x$- and $y$-directions, respectively. The correlation coefficients between $P$ and $I$, and between $P$ and between $I$ at different frequencies are given in Table 4.2. Both the observed correlation coefficients at the maximum resolution of about $1'$, as well as those of the $\mathcal{F}$ resolution data are given. The high correlations between $I$ in different frequency bands are expected if the visible structure in $I$ is mostly due to faint sources. The $C(I, I) < 1$ is due to uncorrelated noise. However, the correlation between $P$ and $I$ is very low, so the
Figure 4.3: Five maps of a part of the Auriga region at frequencies 341, 349, 355, 360 and 375 MHz. The grey scale is polarized intensity $P$, oversampled 5 times, where white is high intensity. Maximum $P$ (in the 375 MHz band) is approximately 33 K. Superimposed are polarization pseudo-vectors, where each line denotes an independent synthesized beam.

Structure in polarization is not due to small-scale variations in synchrotron emission. Instead, the variations in polarized intensity and polarization angle must be due to a combination of two processes: Faraday rotation and depolarization. We will return to these in Sects. 4.4 and 4.5.
Figure 4.4: Two typical parts of $3 \times 5$ beams in the observed field, represented in small graphs of polarization angle $\phi$ against $\lambda^2$, one graph per independent synthesized beam. The left $3 \times 5$ graphs are taken from a part of the field with high polarized intensity ($\sigma > 15$ per frequency band), the right part from a region of lower intensity ($\sigma \approx 5$). The values within parentheses above each graph are $RM$ and reduced $\chi^2$, respectively. In general, $RMs$ are more regular and alike over several beams in high polarized intensity regions, but there is coherence in $RM$ in lower $P$ regions as well.

4.3.3 Polarization angle

In Fig. 4.3 we show, for a $3^\circ \times 2^\circ$ subfield slightly northwest of the center of the mosaic, maps of polarized intensity $P$ (grey scale) and polarization angle $\phi$ (superimposed pseudo-vectors) in the 5 frequency bands. In the grey scale, which has 5 samples per linear beam width, white corresponds to high $P$. The length of the vectors (one per beam) scales with polarized intensity. The amount of structure in polarization angle is highly variable. At high polarized intensities, the polarization angles mostly vary quite smoothly. However, there are also abrupt changes on the scale of a beam, the most conspicuous of which give rise to the depolarization canals.

4.4 Faraday rotation

4.4.1 The derivation of rotation measures

The rotation measure ($RM$) of the Faraday-rotating material follows from $\phi(\lambda^2) \propto RM \lambda^2$. As $\phi = 0.5 \arctan(U/Q)$, the values of polarization angle are ambiguous over $\pm n 180^\circ$ ($n = 1, 2, 3, \ldots$). For most beams, it was possible to obtain a linear $\phi(\lambda^2)$-relation by using angles with a minimum angle difference between neighboring frequency bands. Only in about 1% of the data, addition or subtraction $180^\circ$ to the
polarization angle in at least one of the frequency bands was needed. One could increase $R Ms$ by adding or subtracting multiples of $180^\circ$, but that would lead to $R Ms$ in excess of $100 \text{ rad m}^{-2}$. If such high $R Ms$ were real, bandwidth depolarization (see Section 4.5) would have totally annulled any polarized signal. In addition, the $R Ms$ we obtain using the criterion of minimum angle differences (of order of $10 \text{ rad m}^{-2}$) were also obtained in other observations in this direction (Bingham and Shakeshaft 1967, Spolsttra 1984). So we conclude there are hardly any $180^\circ$-ambiguities in polarization angle and we use the minimum angle difference between frequency bands to determine $R Ms$.

In Fig. 4.4 we show $\phi(\lambda^2)$-plots for small arrays of contiguous, independent beams in typical subfields of the $\delta$-resolution map. In parentheses are the values of the $R M$ and the reduced $\chi^2$ of the linear $\phi(\lambda^2)$-fit. The leftmost 3×5 plots are from a region with high $P$ ($\sigma \geq 15$ in all frequency bands), while the rightmost plots are from a region with lower $P$ ($\sigma \approx 5$). In general, the $\phi(\lambda^2)$-relation is closer to linear in high $P$ regions than in lower $P$ regions, but in the larger part of the Auriga field $R Ms$ are very well-determined.

4.4.2 The distribution of rotation measures

The $R Ms$ in the Auriga-field are plotted as circles in Fig. 4.5. Filled circles denote positive values, open circles negative, and the diameters of the circles are proportional to $R M$. For clarity, we only show the $R Ms$ for one in four independent beams. In addition, $R Ms$ are only shown when they were reliably determined, i.e. for the average polarized intensity $\langle P \rangle > 5\sigma$, and for reduced $\chi^2$ of the fit $< 2$. The left panel shows $R M$ as determined from the observations. Most $R Ms$ are negative; the few positive values occur for $\delta \leq 52^\circ$, while the largest negative values occur for $\delta \geq 55^\circ$. We modeled this systematic variation by a large-scale linear gradient, and find a gradient of about 1 rad m$^{-2}$ per degree, with the steepest slope along position angle (from north through east) of $-20^\circ$. The average $R M$, i.e. the $R M$ value in the center of the fitted $R M$-plane, over the field is $R M_0 \approx -3.4 \text{ rad m}^{-2}$. Subtracting the best-fit gradient from the $R M$ distribution yields the $R Ms$ shown in the right-hand panel in Fig. 4.5 with only small-scale structure in $R M$. The distribution of small-scale $R Ms$ (i.e. $R M$ where the large-scale gradient is subtracted) is significantly narrower than the total $R M$ distribution ($\sigma_{R M} = 1.4 \text{ K}$ instead of $2.3 \text{ K}$), and is more symmetrical, as is shown in Fig. 4.6, where the two histograms are compared.

The decomposition of the $R M$-distribution into a constant component, a gradient, and a small-scale component is not physical, as there is probably structure on all scales. However, for our purpose it is a good approximation, and it allows us to estimate the large-scale and random components of the Galactic magnetic field, see Section 4.7.

4.4.3 Structure in $R M$ on arcminute scales

Although the $R M$-distribution shows structure on many scales, the most intriguing changes occur on the scale of the beam, as illustrated in Fig. 4.7. In the figure, we display a 9×7 array of $\phi(\lambda^2)$-plots, overlaid on a grey scale representation of $P$ at 349 MHz. Although $R M$ varies quite smoothly over most of the area, there are also
**Figure 4.5:** Rotation measures given as circles with a diameter proportional to the magnitude of the RM, superimposed over $P$ at 349 MHz in grey scale. The scaling is in rad m$^{-2}$. RMs are only shown if $P > 5\sigma$ and reduced $\chi^2 < 2$, and in both directions, only every second independent beam is plotted for clarity. The right map shows the same data where the best-fitting linear gradient in RM is subtracted.

**Figure 4.6:** Histogram of RMs in the Auriga field. The dashed line gives the distribution of all “reliably determined” RMs, i.e. where $P > 5\sigma$ and reduced $\chi^2 < 2$. The solid line gives the reliably determined RM distribution after subtracting a best-fit linear RM gradient from the data. Dotted lines show Gaussian fits to the two histograms.
Figure 4.7: White symbols and lines are plots of $\phi$ against $\lambda^2$ and their linear fits, one for each independent beam, so that the slope is $RM$. Sudden $RM$ changes occur over one beam width. The grey scale is polarized intensity at 349 MHz oversampled by a factor of 5. Maximum $P$ (lightest grey) is $\sim 40$ mJy/beam, while in the bottom left pixel $P$ is set to zero for comparison.

Abrupt changes, which frequently involve a change of the sign of $RM$ (e.g. at $(\alpha, \delta) = (94.60, 53.00)^\circ$ or $(94.68, 53.20)^\circ$). Abrupt $RM$ changes over one beam with the right magnitude to cause an angle change of $\pm(n + 1/2)180^\circ$ cause depolarization canals. As $RM$ is an integral over the entire line of sight, it is difficult to understand such changes in $RM$ as resulting from structure in the magnetic field and/or distribution of electrons. As a matter of fact, numerical simulations of the warm ISM show that beam depolarization can enhance $RM$ differences and greatly steeper the gradient (Chapter 9). In addition, models of a medium that produces Faraday rotation and that emits synchrotron radiation show that the $RM$ can change sign without a corresponding change in the direction of the magnetic field (Sokoloff et al. 1998, Chi et al. 1997).

4.5 The structure in $P$, and the implied properties of the ISM

As $I$ and $P$ are uncorrelated, the structure in $P$ cannot be caused by structure in emission, but instead is created by several instrumental and/or depolarization effects. We discuss these processes in detail in Chapter 3. Here we summarize them, and discuss the results for the Auriga region.

Structure in $P$ could be induced by the insensitivity of the WSRT to large-scale
structure, as a result of missing short-spacing information. However, it is very unlikely that this effect is important in the Auriga observations, because the distribution of $RMs$ per pointing center is too wide to allow offsets, as explained in Chapter 3. If the distribution of $RMs$ in a Faraday screen is too wide, the distributions of $Q$ and $U$ of a polarized background that is Faraday-rotated by the screen have averages very close to zero (even if the background was uniformly polarized). As a result, the missing short-spacing information does not produce offsets. For our frequencies the latter is true if $\sigma_{RM} \geq 1.8 \text{ rad m}^{-2}$. In the Auriga region, this condition is met in essentially all of the subfields of the mosaic. Furthermore, a test in which offsets are added (described in the same Chapter) to produce better linearity of the $\phi(\lambda^2)$-relation, did not result in reliable non-zero offsets.

Structure in $P$ could also be generated by non-constant, but high, $RM$ values as a result of bandwidth depolarization. Summation of polarization vectors with different angles in a frequency band will reduce the polarized intensity. The amount of depolarization depends on the spread in polarization angle within the band which, in turn, depends on the $RM$ and the width of the frequency band. For the $RMs$ of order 10 rad m$^{-2}$ and our bandwidth of 5 MHz the effect is negligible.

A special kind of structure in $P$ is due to beam depolarization, viz. the cancellation of polarization by vector summation across the beam defined by the instrument. This is a 2-dimensional summation of polarization vectors across the plane of the sky. This process is probably responsible for the depolarization canals described by Haverkorn et al. (2000, Chapter 6). It certainly cannot produce all of the structure in $P$ because it occurs on a single, small scale, whereas $P$ has structure on a large range of scales.

Finally, depth depolarization occurs along the line of sight in a medium that both emits synchrotron radiation and produces Faraday rotation. It must be responsible for a large part of the structure in $P$. The ISM in the Galactic disk most likely is such a medium. The Galactic synchrotron emission has been modeled by Beermann et al. (1985) as the sum of two contributions: the thin disk with a half equivalent width of $\sim 180$ pc, which coincides with the stellar disk, and a thick disk of half equivalent width $\sim 1800$ pc (where the galactocentric radius of the sun is assumed $R_\odot = 10$ kpc). Thermal electrons exist also out to a scale height of approximately a kpc, contained in the Reynolds layer (Reynolds 1989, 1991). So up to a height of about a kpc, the medium probably consists of both emitting and Faraday-rotating material.

In such a medium, the polarization plane of the radiation emitted at different depths is Faraday-rotated by different amounts along the line of sight. Therefore the polarization vectors cancel partly, causing depolarization along the line of sight even if the magnetic field and thermal and relativistic electron densities are constant (differential Faraday rotation, see Gardner and Whiteoak 1966, Burn 1966, Sokoloff et al. 1998). If the medium contains a random magnetic field, the polarization plane of the synchrotron emission will vary along the line of sight causing extra depolarization in addition to the differential Faraday rotation (called internal Faraday dispersion). We will refer to the combination of the two mechanisms along one line of sight as depth depolarization, thus adopting the terminology proposed by Landecker et al. (2001).

Depth depolarization is the dominant depolarization process in the Auriga field. We constructed a numerical model in Chapter 3 to evaluate if depth depolarization can indeed be held responsible for most of the structure in $P$, in the Auriga region as well.
as in the other region that we studied. At the same time, the model makes it possible to estimate the conditions in the warm component of the ISM. The model consists of a layer corresponding to the thin Galactic disk, where random magnetic fields on small scales are present, and a constant background polarized radiation.

In the model, the thin disk (with height $h$) is divided in cells of a single size $d$, which contain a random magnetic field component $B_{\text{ran}}$ and a regular magnetic field $B_{\text{reg}}$. The amplitude of the random magnetic field $B_{\text{ran}}$ is constant, but in each cell its direction is randomly chosen. Each cell emits synchrotron radiation, proportional to $B_0^2 = (B_{\text{ran}, \perp})^2 + (B_{\text{reg}, \perp})^2$, where we add energy densities because the two components are physically separated different fields. Only a fraction $f$ of the cells contains thermal electrons, and therefore Faraday rotates all incoming radiation (to represent the filling factor $f$ of the warm ISM).

We distinguish several kinds of model parameters. First, we have input parameters with values taken from the literature; these are thermal electron density $n_e = 0.08 \text{ cm}^{-3}$, filling factor $f = 0.2$ (both from Reynolds 1991), and height of the Galactic thin disk $h = 180$ pc (Beermann et al. 1985). Secondly, there are constraints from the observations of the Auriga field, viz. the width, shape and mean of the $Q$, $U$, $I$ and $R$ distributions. In addition, we define parameters without observational constraints, such as cell size $d$ and emissivity per cell, which we vary in a attempt to constrain them from the model given the observational constraints. Finally, there are parameters that can be adjusted and optimized for each allowed cell size and emissivity per cell, like the random and regular magnetic field components and background intensity.

Using the observations of the Auriga region to constrain the parameters in the depth depolarization model, we obtain the following estimates. The cell size $d$ is probably in the range of 5 - 20 pc, the random component of the magnetic field $B_{\text{ran}}$ is about 1 $\mu$G for a cell size of 5 pc or larger, and up to 4 $\mu$G for smaller cell sizes. With the estimate of the large-scale, regular field ($\sim 3$ $\mu$G) this leads to a value of about 0.3 to 1.3 for $B_{\text{ran}}/B_{\text{reg}}$. This is lower than most other estimates of $B_{\text{ran}}/B_{\text{reg}}$ in the literature, which seem to indicate a value larger than 1. One possible explanation for this discrepancy is the fact that the Auriga region is not very large. Therefore random magnetic field components that project to angular scales larger than a few degrees will have been interpreted as components of the regular field. If such large-scale random components exist, this effect will artificially decrease $B_{\text{ran}}/B_{\text{reg}}$. Secondly, the Auriga field was selected for its conspicuous polarization structure, and high polarized intensity in general indicates a more regular magnetic field structure. Finally, the Auriga region is observed in an inter-arm region. The next spiral arm, the Perseus arm, is more than 2 kpc away, and we expect that at these frequencies, all polarized emission from the Perseus arm is depolarized. There are indications that in the inter-arm region, regular magnetic field component is higher than the random magnetic field component (Han and Qiao 1994, Indrani and Deshpande 1998, Beck 2001).

A cell of size 10 pc located at the far edge of the thin disk still subtends a little less than a degree in our maps. We see structure on smaller scales, so that cells of approximately a parsec have to be present as well. Indeed, there are many indications that structure exists in the ISM on many scales (Armstrong et al. 1995, Minter and Spangler 1996).

The distance out to which polarized radiation is not depolarized by depth depolar-
ization, as derived from the model, is shown in Fig. 4.8. The total path length through the medium containing small-scale structure was about 600 pc, which corresponds to a height of the thin disk of about 150 pc for the Auriga latitude of $\sim 16^\circ$. Radiation emitted at large distances is depolarized more than emission from nearby, but there is no definite polarization horizon due to depth depolarization. Beam depolarization can cause a polarization horizon, as structure further away will be on smaller angular scales, which causes more beam depolarization (Landecker et al. 2001).

### 4.6 Polarized extragalactic point sources

Thirteen polarized extragalactic point sources in the Auriga-field were selected based on three selection criteria: (1) signal-to-noise $\sigma > 3$ in $P$ at all frequencies, (2) the degree of polarization $p$ is higher than 1% at all frequencies (because a lower polarization degree can also be caused by instrumental polarization, see Section 4.2), and (3) the source does not lie too far at the edge of the field away from any pointing center, to limit the contribution from instrumental polarization. The 13 sources are all detected in the $1\prime$ resolution data, and most of them are unresolved at that resolution. Their $RMs$ vary from approximately $-13$ to 13 rad m$^{-2}$.

The relevant properties of the sources are given in Table 4.3, where the second column gives the right ascension and declination, column 3 the $RM$, and column 4 the reduced $\chi^2$ of the $\phi(\lambda^2)$-fits. Columns 5, 6, and 7 give the values of $P, I$ and percentage of polarization $p$ respectively, averaged over all frequency bands. Fig. 4.9 shows the polarization angle $\phi$ plotted against $\lambda^2$ for each point source.

In Fig. 4.10, we show the 13 sources overlaid on a grey scale plot of polarized intensity at 349 MHz, with their $RMs$ indicated. The diameter of the symbol and its shading indicate the value of $RM$.

Sources 1 and 2 could have a significant contribution of instrumental polarization, because they are observed in only one pointing center and are located about 0.8$^\circ$ to 1$^\circ$ away from the pointing center. Nevertheless, they show linear $\phi(\lambda^2)$-relations, which
Table 4.3: Extragalactic sources in the Auriga region. The second column gives position, and the third \( RMs \). Reduced \( \chi^2 \) of the \( \phi(\lambda^2) \) relation is given in column 4. Columns 5, 6 and 7 give resp. \( P, I \) (both in mJy/beam) and percentage of polarization \( p \) averaged over the five frequency bands.

<table>
<thead>
<tr>
<th>No.</th>
<th>((\alpha, \delta) \ (^{\circ}, {^\circ}))</th>
<th>(RM ) (rad m(^{-2}))</th>
<th>(\chi^2)</th>
<th>(\langle P \rangle)</th>
<th>(\langle I \rangle)</th>
<th>(\langle p \rangle)</th>
</tr>
</thead>
<tbody>
<tr>
<td>1</td>
<td>[97.86, 55.32]</td>
<td>7.4 (\pm) 0.2</td>
<td>12.9</td>
<td>456</td>
<td>2021</td>
<td>2.3</td>
</tr>
<tr>
<td>2</td>
<td>[97.74, 53.61]</td>
<td>12.9 (\pm) 0.3</td>
<td>3.1</td>
<td>149</td>
<td>216</td>
<td>6.9</td>
</tr>
<tr>
<td>3</td>
<td>[96.61, 53.32]</td>
<td>6.6 (\pm) 0.4</td>
<td>1.7</td>
<td>150</td>
<td>224</td>
<td>6.7</td>
</tr>
<tr>
<td>4</td>
<td>[96.49, 54.07]</td>
<td>5.7 (\pm) 0.4</td>
<td>3.7</td>
<td>119</td>
<td>380</td>
<td>3.1</td>
</tr>
<tr>
<td>5</td>
<td>[95.31, 50.75]</td>
<td>0.6 (\pm) 0.4</td>
<td>1.6</td>
<td>147</td>
<td>165</td>
<td>8.8</td>
</tr>
<tr>
<td>6</td>
<td>[95.21, 50.75]</td>
<td>3.1 (\pm) 0.4</td>
<td>0.2</td>
<td>14.7</td>
<td>197</td>
<td>7.5</td>
</tr>
<tr>
<td>7</td>
<td>[94.72, 54.80]</td>
<td>7.9 (\pm) 0.3</td>
<td>3.0</td>
<td>18.8</td>
<td>282</td>
<td>6.7</td>
</tr>
<tr>
<td>8</td>
<td>[93.39, 55.61]</td>
<td>6.3 (\pm) 0.2</td>
<td>1.1</td>
<td>21.3</td>
<td>180</td>
<td>11.8</td>
</tr>
<tr>
<td>9</td>
<td>[92.32, 52.92]</td>
<td>2.7 (\pm) 0.1</td>
<td>6.1</td>
<td>35.8</td>
<td>1222</td>
<td>2.9</td>
</tr>
<tr>
<td>10</td>
<td>[90.25, 50.37]</td>
<td>(-10.2 \pm 1.4)</td>
<td>4.4</td>
<td>7.4</td>
<td>168</td>
<td>4.4</td>
</tr>
<tr>
<td>11</td>
<td>[89.31, 54.98]</td>
<td>3.0 (\pm) 0.4</td>
<td>1.4</td>
<td>11.6</td>
<td>340</td>
<td>3.4</td>
</tr>
<tr>
<td>12</td>
<td>[89.20, 52.50]</td>
<td>(-8.1 \pm 0.1)</td>
<td>12.7</td>
<td>22.1</td>
<td>477</td>
<td>4.6</td>
</tr>
<tr>
<td>13</td>
<td>[88.95, 52.27]</td>
<td>(-13.6 \pm 0.4)</td>
<td>5.8</td>
<td>14</td>
<td>237</td>
<td>5.9</td>
</tr>
</tbody>
</table>

Figure 4.9: Plots of polarization angle \( \phi \) against \( \lambda^2 \) for the 13 polarized extragalactic sources in Table 4.3. Note that the scaling of the \( \phi \)-axis differs for each source.
indicates that instrumental polarization is not important. These sources are therefore included in the analysis.

The strong correlation of RM across the sky indicates a Galactic component to the RMs of the extragalactic point sources. The best-fitting linear gradient to the RMs of the sources has a steepest slope of 3.62 rad m\(^{-2}\) per degree in position angle 72\(^\circ\), i.e. roughly in the direction of increasing Galactic latitude (position angle 66\(^\circ\)). The standard deviation of RMs around this gradient is \(\sigma_{RM} \approx 3.6\) rad m\(^{-2}\). This is consistent with the result of Leahy (1987), who finds a typical 'internal' RM contribution from the extragalactic source or a halo around it of \(\sim 5\) rad m\(^{-2}\).

The change of sign of RM of the extragalactic sources within the observed region indicates a reversal in the Galactic magnetic field. (Unlike the diffuse emission, sign changes in RMs of extragalactic sources cannot be due to depolarization effects.) The reversal exists on scales larger than our field of view, as is shown in Fig. 4.11, where we combine the RMs of our sources with those from the literature. The circles are our sources, the squares indicate sources that were detected by Simard-Normandin et al. (1981) and/or by Tabara and Inoue (1980), and the one triangle indicates the only pulsar nearby (Hamilton and Lynne 1987). The numbers next to the squares and triangle give the magnitude of RM of the source (the diameter of the symbol is only proportional to RM up to \(RM = 15\) rad m\(^{-2}\)). One source at \((\alpha, \delta) = (91.25^\circ, 48^\circ)\) was omitted, because Simard-Normandin et al. give \(RM = 34\) rad m\(^{-2}\) and Tabara and Inoue give \(RM = -64.9\) rad m\(^{-2}\).

The source distribution shows a clear magnetic field reversal, which is not on Galactic scale but on smaller scales, as is shown in Figs. 1 and 2 of Simard-Normandin and
Figure 4.11: RMs of extragalactic point sources and of one pulsar. The circles are the point sources detected in the Auriga field, where the radius of the circle is proportional to the magnitude of $R.M$. Squares are extragalactic point sources detected by Simard-Normandin et al. (1981) and/or by Tabara and Inone (1980), and the one triangle in the field denotes the only pulsar near the Auriga field (Hamilton and Lyne 1987). Values of the $R.M$ are written next to the sources, and for $R.M < 15$ rad m$^{-2}$ the radius of the symbol is proportional to $R.M$.

Kronberg (1980), which show RMs of extragalactic sources over the whole sky.

The gradient in $R.M$ measured in the extragalactic point sources is completely unrelated and almost perpendicular to the structure in $R.M$ from the diffuse emission. This can be explained from the widely different path lengths that extragalactic sources and diffuse emission probe. Diffuse emission can only be observed out to a distance of a few hundred pc to a kpc, as radiation from further away will be mostly depolarized. On the contrary, extragalactic sources are Faraday-rotated over the complete path length through the Milky Way of many kiloparsecs long. Therefore, the $R.M$ structure from diffuse radiation and from extragalactic sources gives information about different regions in the Galaxy.

4.7 Discussion

4.7.1 High $P$-structures and alignment with the Galactic plane

The depth depolarization model described in Section 4.5 only describes the properties of the ISM in a statistical manner, by giving a representation of an average line of sight. Therefore, it does not pretend to give a detailed description of the distribution of $P$, as
it is actually observed. Even the global properties of the $P$ distribution, like the general alignment with the Galactic plane, is not part of the model. This alignment is not only visible in the Auriga field, but also in other regions observed with the WSRT, such as that centered on $(l, b) = (137^\circ, 7^\circ)$ (Chapter 5) and those for which we have single-frequency polarization data from the WENSS (Chapter 7). The large-scale structures in $P$ cannot be caused by substantial extra emission, because that should be accompanied by corresponding structure in $I$, which is not observed. Instead, where the polarized intensity is highest, the depolarization is probably the least, so that there the ISM is relatively transparent to polarized emission. In general, high depolarization is caused by a large amount of structure in $RM$, along the line of sight and/or over the sky (the latter on the scale of the beam or smaller). This explanation of the structure of high $P$ is borne out by the relation between polarized intensity $P$ and $\sigma_{RM}$, shown in Fig. 4.12. This figure shows the width $\sigma_{RM}$ of a Gaussian fit to the $RM$ distribution in several intervals of polarized intensity with a width $\Delta P = 0.01$ Jy/beam. Rotation measure clearly varies more at lower polarized intensity, where only well-determined $RMs$ are used so that the effect cannot be due to noise. For the two intervals of highest $P$ ($0.08 - 0.09$ Jy/beam and $0.09 - 0.10$ Jy/beam), not shown in Fig. 4.12, no Gaussian fit could be made, as the distributions were bimodal. The peaks in these bimodal distributions correspond to two spatially coherent structures, which are shown in the two panels of Fig. 4.13. In the region around $(\alpha, \delta) = (94.6^\circ, 51.8^\circ)$, $-1.5 \leq RM \leq 0$ rad m$^{-2}$, while in the region around $(\alpha, \delta) = (92.2^\circ, 53.8^\circ)$, $0 \leq RM \leq 1.3$ rad m$^{-2}$. Clearly, $RM$ is indeed very uniform in regions of high polarization. The low variation in $RM$ is also reflected in the uniformity of the polarization angles.

A constant $RM$ over a certain area sets an upper limit on structure in magnetic field and thermal electron density. The fact that some of the filamentary structure in $P$ is aligned with Galactic latitude suggests a magnetic origin, although stratification of $n_e$ over Galactic latitude can contribute as well. For variation of polarization angle over the filament of less than about $20^\circ$ (from Fig. 4.13), $\Delta RM \leq 0.5$ rad m$^{-2}$ is needed. If magnetic field structure would be the same over the entire path length
towards the filament, and assuming \( n_e = 0.08 \text{ cm}^{-3} \), this requires a magnetic field change \( \Delta B_\| \leq 0.01 \mu \text{G} \) with a polarization depth of 600 pc. Even if the filamentary structures are sheet-like with a constant magnetic field over a large part of the path length, or if the polarization horizon is locally much shorter than the highly uncertain value of 600 pc, the upper limit to \( \Delta B \) is very low. Therefore, this structure of high \( P \) aligned with Galactic latitude indicates the existence of long filaments or sheets parallel to the Galactic plane of highly uniform magnetic field and thermal electron density. As back- and/or foreground structure can also imprint structure in \( RM \), these regions may well be characterized by low thermal electron density.

4.7.2 The Galactic magnetic field

The Auriga field of \( 7^\circ \times 9^\circ \) represents only 0.1% of the sky, and is therefore by itself not very well suited for an analysis of the global properties of the Galactic magnetic field. However, we can estimate strength and direction of the regular component of the Galactic magnetic field from the distribution of \( RM \)s of the diffuse radiation and of the extragalactic sources over the field.

First, we need to determine how much of the \( RM \) gradient seen in the diffuse emission is caused by \( n_e \), and how much by \( B_\| \). Independent measurements of \( n_e \) can be obtained from the emission measure \( EM = \int n_e^2 \, dl \), as measured with the Wisconsin H\( \alpha \) Mapper survey (WHAM, Haffner et al., in prep, Reynolds et al. 1998), shown in Fig. 4.14. As the resolution is about a degree, the WHAM survey cannot be used to determine small-scale structure in \( n_e \), but gives an estimate of the global gradient in
Figure 4.14: Superimposed circles are emission measures from the WHAM Hα survey, to a maximum intensity of ~5 Rayleigh, at a resolution of one degree. The grey scale denotes $P$ at 349 MHz.

$n_e$ over the field. From north to south, the Hα observations show an increase from about 1 to 5 Rayleighs (1 Rayleigh is equal to a brightness of $10^6/4\pi$ photons cm$^{-2}$ s$^{-1}$ ster$^{-1}$, and corresponds to an $E M$ of about 2 cm$^{-6}$ pc for gas with a temperature $T = 10000$ K). The small enhancement in $E M$ at $(\alpha, \delta) = (94.8, 55.5)$ is probably related to the ancient planetary nebula PnWe1 (PW1, Tweedy and Kwitter 1996). Note that the Hα emission probes a longer line of sight than the diffuse polarization, and is velocity integrated over ~80 to 80 km s$^{-1}$ LSR, so complete correspondence is not expected. However, a gradient in $RM$ that was solely due to the $n_e$-gradient would have a sign opposite to that of the observed $RM$-gradient. Therefore, the observed gradient in $RM$ must be due to a gradient in the parallel component of the magnetic field. If the gradient in Hα originates in the same medium as the polarized radiation, the gradient in magnetic field must even be large enough to compensate the gradient in electron density in the opposite direction.

So the $RM$ gradient of the diffuse polarized emission of about 1 rad m$^{-2}$ per degree in the direction of positive Galactic longitude must be caused by a corresponding increase of the parallel component of the Galactic magnetic field. The direction of the gradient in longitude and the sign of the average $RM$ are consistent with a regular Galactic
magnetic field directed away from us in the second quadrant, as found from other observations (see e.g. review by Han, 2001). Attributing the observed $RM$-gradient completely to the change in the parallel component of the Galactic magnetic field, we can derive the strength of its regular component. The average $RM_0$ and the value of the $RM$ gradient give two independent estimates of the regular magnetic field, if a certain pitch angle is assumed. Assuming a path length of 600 pc and electron density $n_e = 0.08 \text{ cm}^{-3}$, the two magnetic field determinations agree on a regular magnetic field of $B_{eg} \approx 1 \mu G$ for a pitch angle $p = -14^\circ$, in agreement with earlier estimates (Vallée 1995).

The gradient in the $RMs$ of the extragalactic sources does not show a longitudinal component at all, contrary to that in the $RMs$ of the diffuse polarization observations. This indicates that other large-scale magnetic fields become important, fields that vary on scales larger than our field. Observations of halos of external galaxies (e.g., Dumke et al. 1995) have shown that cell sizes in the Galactic halo are much larger than in the disk, from 100 - 1000 pc. Even at a distance of 3000 pc, the size of the total Auriga field is $\sim 400$ pc, so big cells in the halo could be responsible for a considerable part of what we call the regular component of the magnetic field, as measured in the $RMs$ of the extragalactic sources.

We subtracted the gradient of the diffuse emission from the gradient in the $RMs$ of the extragalactic sources, after which a gradient in $RM$ of $\Delta RM \approx 3.6 \text{ rad m}^{-2}$ in position angle $54^\circ$ remains (only deviating by about $10^\circ$ from the direction perpendicular to the Galactic plane). This indicates a regular magnetic field of about $1 \mu G$ in this direction, assuming that $n_e$ and $B$ remain constant over a total line of sight of 3000 pc. The magnetic field strengths actually present in the medium will be higher if the field varies along the line of sight.

4.8 Conclusions

Multi-frequency polarization observations of the diffuse Galactic background yield information on structure in the Galactic warm gas and magnetic field.

The multi-frequency WSRT observations of a region of size $\sim 7^\circ \times 9^\circ$ at $l = 161^\circ$, $b = 16^\circ$ in the constellation Auriga show a smooth total intensity $I$, but abundant structure in Stokes parameters $Q$ and $U$ on several scales, with a maximum $T_{b, pol} \approx 13 \text{ K}$ (about 30% polarization). Filamentary structure up to many degrees long is present in $P$, sometimes aligned with Galactic latitude. In addition, narrow, one-beam wide depolarization canals are most likely created by beam depolarization and may well indicate abrupt $RM$ changes. As the polarization structure is uncorrelated to $I$, the structure in $P$ cannot be created by variations in synchrotron emission, but has to be due to Faraday rotation and depolarization mechanisms.

Rotation measure maps show abundant structure on many scales, including a linear gradient of $\sim 1 \text{ rad m}^{-2}$ per degree in the direction of positive Galactic longitude and an average $RM_0 \approx -3.4 \text{ rad m}^{-2}$. The gradient is consistent with the regular Galactic magnetic field if its strength is $B_{eg} \approx 1 \mu G$ and the field is azimuthally oriented with a pitch angle $p \approx -14^\circ$. Ubiquitous structure is present in the $RM$ map on beam size scales ($\theta$), indicating significant changes in magnetic field and/or thermal electron
density over very small spatial scales, although some of the structure may be created by beam depolarization.

There are two dominant depolarization mechanisms which create structure in $P$ in the Auriga field. First, beam depolarization (i.e. depolarization due to averaging out polarization angle structure within one beam width) most likely creates the canals, and is important in regions of low $P$. Additional depolarization is needed to explain the observed $P$ distribution, which is only possible if the medium both Faraday-rotates and emits synchrotron emission. Then, emission originating in the medium can be depolarized along its path towards the observer, so-called depth depolarization. The polarization angle along the path can vary due to Faraday rotation, or due to varying intrinsic polarization angle of the emission. This indicates a varying magnetic field, which however is constrained by the upper limit on structure in $I$. A depth depolarization model was constructed of a layer of cells with varying magnetic fields and a constant background polarization. Constraints from the Auriga observations give estimates for several parameters in the ISM. The cell size of structure in the ISM is constrained to $\sim 15$ pc, and the ratio of random to regular field is $0.7 \pm 0.5$.

Thirteen extragalactic sources in the Auriga field also show a gradient in $R M$, but roughly in the direction of Galactic latitude, which is perpendicular to the gradient in the diffuse emission. The $R M$ distributions from diffuse radiation and from extragalactic point sources are so different because the diffuse radiation mostly probes the first few hundred parsecs, whereas $R M$s from the point sources are built up along the entire line of sight through the Milky Way. $R M$s of extragalactic sources change sign over the field, which indicates a local reversal of the magnetic field.

Acknowledgements

We thank R. Beck, E. Berkhuijsen and F. Heitsch for critical reading and useful comments. The Westerbork Synthesis Radio Telescope is operated by the Netherlands Foundation for Research in Astronomy (ASTRON) with financial support from the Netherlands Organization for Scientific Research (NWO). The Wisconsin H-Alpha Mapper is funded by the National Science Foundation. MH is supported by NWO grant 614-21-006.

References

Beck R., 2001, SSRv 99, 243
Bridle A. H., Davis M. M., Foncalent E. B., & Loqueux J., 1972, NPhS 235, 123
Ferrière K. M., 2001, RvMP 73, 1031
Gardner F. F., & Whiteoak J. B., 1966, ARAA 4, 245
Indrani C., & Deshpande A. A., 1998, NewA 4, 331
Landecker T. L., Uyaniker B., & Kothes R., 2001, AAS 199, #58.07
Simard-Normandin M., & Kronberg P. P., 1979, Nature 270, 115
Multi-frequency polarimetry of the Galactic radio background around 350 MHz: II. A region in Horologium around $l = 137^\circ$, $b = 7^\circ$

M. Haverkorn, P. Katgert and A. G. de Bruyn, submitted to A&A

Abstract

A conspicuous ring-like structure with a radius of about $1.4^\circ$ was studied with the Westerbork Synthesis Radio Telescope (WSRT) at 5 frequencies around 350 MHz. This ring is very prominent in Stokes $Q$ and $U$ and in polarization angle, and less so in polarized intensity $P$. No corresponding structure is visible in total intensity Stokes $I$, which indicates that the ring is created by Faraday rotation and depolarization processes. The polarization angle changes from the center of the ring outwards to a radius $> 1.7^\circ$. Thus, the structure in polarization angle is not ring-like but resembles a disk, and it is larger than the ring in $P$. The Rotation measure ($RM$) decreases almost continuously over the disk, from $RM \approx 0$ rad m$^{-1}$ at the edge, to $-8$ rad m$^{-1}$ in the center, while outside the ring, the $RM$ is slightly positive. This radial variation of $RM$ yields stringent constraints on the nature of the ring-like structure, because it rules out any spherically symmetrical magnetic field configuration, such as might be expected from supernova remnants or wind-blown bubbles. We discuss several possible connections between the ring and known objects in the ISM, and conclude that the ring is a predominantly magnetic funnel-like structure. This description can explain both the field reversal from outside to inside the ring, and the increase in magnetic field, probably combined with an electron density increase, towards the center of the ring. The ring-structure in $P$ is most likely caused by a lack of depolarization due to a very uniform $RM$ distribution at that radius. Beyond the ring, the $RM$ gradient increases, depolarizing the polarized emission, so that the polarized intensity decreases. In the southwestern corner of the field a pattern of narrow filaments of low polarization, aligned with Galactic longitude, is observed, indicative of beam depolarization due to abrupt changes in $RM$. This explanation is supported by the observed $RM$. 

87
5.1 Introduction

Observation of the diffuse polarized background of our Galaxy at radio frequencies provides a unique way to observe features in the ISM that are otherwise invisible. In this paper we report and analyze the detection in radio polarization of a remarkable circular structure of $\sim 3^\circ$ in size.

The observed radiation is synchrotron radiation, originating from relativistic cosmic ray electrons in interaction with the Galactic magnetic field. The linearly polarized component of the synchrotron radiation is modulated by several mechanisms, among which Faraday rotation, viz. the birefringence of left and right handed circularly polarized radiation in a medium which contains a magnetic field and free electrons. For linear polarization, this results in a rotation of the angle of polarization $\phi$ in passage through a magneto-ionized medium. This rotation is proportional to the square of the wavelength where the proportionality constant is the rotation measure ($RM$). The $RM$ depends on the magnetic field component parallel to the line of sight $B_\parallel$, weighted by the thermal electron density $n_e$, integrated over the line of sight. Thus multi-wavelength observations of linear polarization yield directly the electron-density-weighted value of the interstellar magnetic field.

The circular structure we have observed at $(l, b) \approx (137^\circ, 7^\circ)$ is located in the middle of a region of very high polarization extending over many degrees as measured by e.g. Brouw and Spoestra (1976).

Bingham and Shakeshaft (1967) were the first to observe the circular structure in this region in a map of $RM$ that they constructed from surveys by Berkhuijsen et al. (1963, 1964) at 408 MHz and 610 MHz, by Wielebinski and Shakeshaft (1964) at 408 MHz and by Bingham (1966) at 1407 MHz. Their $RM$ map shows a circular structure of about the size of their beam ($\approx 2^\circ$) where $RM \approx -5$ rad m$^{-2}$, in a region where $RMs$ are $-1$ rad m$^{-2}$ $< RM < 0$ rad m$^{-2}$. Verschuur (1968) reobserved this region at higher angular resolution ($\approx 40''$) with the 250ft MkI telescope at Jodrell Bank at 408 MHz. Verschuur found that it is a ring-like structure of low polarized intensity (50% less than in its surroundings) with a radius of a few degrees, and with a small region (about one beam) of almost zero polarization at the southwestern edge of the ring. Verschuur suggests that the region is connected to a bright B2Ve star HD20336 which is located close to the center of the ring.

In a later paper, Verschuur (1969) presents HI measurements obtained with the 300ft Green Bank telescope which show an deficiency in HI along the trajectory of the star HD20336. He explains this by assuming the star is moving through the neutral medium, expelling the HI and ionizing the remaining small fraction of neutral material. He does not mention the ring-like structure.

The ring-like structure, which is mainly visible in Stokes $Q$ and $U$ and in polarization angle, was noticed in polarization maps produced as a by-product of the Westerbork Northern Sky Survey (WENSS, Rengelink et al. 1997, Schnitzeler et al., in prep., Chapter 7). In 1995/1996, the region was reobserved with the Westerbork Synthesis Radio Telescope (WSRT) at 8 frequencies by T. Spoestra, who kindly allowed us to analyze his observations.

In Section 5.2, we discuss the observations, while in Section 5.3 the observational results are given. Section 5.4 presents some observed properties of the ring-like struc-
ture, that are subsequently interpreted in Section 5.5. Possible connections of the ring structure to other observed features in the Galaxy are presented in Section 5.6. In Section 5.7, we discuss remarkable structure in the field unrelated to the ring, and Section 5.8 presents our conclusions.

5.2 The observations

The Westerbork Synthesis Radio Telescope (WSRT) was used to observe a mosaic of approximately $7^\circ \times 7^\circ$, centered around $(l, b) \approx (137^\circ, 7^\circ)$ in the constellation of Horologium. Data were taken in 8 frequency bands simultaneously with a band width of 5 MHz, but 3 frequency bands contain no usable data due to radio interference. The 5 bands which contain good data are centered at 341 MHz, 349 MHz, 355 MHz, 360 MHz, and 375 MHz. The maximum resolution of the WSRT array is $\sim 1\,\text{arcsec}$, but the data were smoothed using a Gaussian taper to obtain a $5.0\,\text{arcsec}$ resolution. The derived maps of linearly polarized intensities Stokes $Q$ and $U$ were used to compute the polarized intensity $P$ and polarization angle $\psi$. The noise in the polarization data was derived from the rms signal in (empty) Stokes $V$ maps at $5.0\,\text{arcsec}$ resolution and was $\sim 5\,\text{mJy/beam}$ for all bands, see Table 5.1.

To obtain a large field of view, and to reduce off-axis instrumental polarization, the mosaic technique was used. The telescope cycled through a grid of pointing positions during the 12hr observation, integrating 50 seconds per pointing, so that every pointing position was observed many times per 12hr period. The mosaic was constructed from $5 \times 5$ pointings of $3^\circ \times 3^\circ$ each, where the distance between the pointing centers is $1.25^\circ$. The edges of the mosaic far away from a pointing center display high instrumental polarization, and these were not used in the analysis (see Chapter 2).

The data was reduced with the NEWSTAR reduction package, using the unpolarized calibrator sources 3C48, 3C147 and 3C286, and the polarized calibrators 3C345 and 3C303. The absolute flux scale at 325 MHz is based on a flux density of 26.93 Jy for 3C286 (Baars et al. 1977). For details on the data reduction procedure, see Chapter 2.

To avoid radio interference from the Sun, the observations were taken mostly or completely at night, as shown in Table 5.1. Because the observations were done in winter, and during a solar minimum, the total electron content (TEC) of the ionosphere is low. These observing conditions minimize the ionospheric Faraday rotation, as is discussed in Chapter 2. We can estimate the $RM$ contribution of the ionosphere from the TEC of the ionosphere and the earth magnetic field. The TEC of the ionosphere above Westerbork at night, in a solar minimum, and in winter is minimal: TEC $\approx 2.2 \times 10^{16} \text{ electrons cm}^{-2}$ (Campbell, private communication). Assuming a vertical component of the earth magnetic field of 4.5 G towards us, and a path length through the ionosphere of 300 km, the $RM$ caused by the ionosphere is $-0.25 \text{ rad m}^{-2}$ at hour angle zero. At larger hour angles, this is even less. So we expect the rotation measure values given in this chapter not to be affected by ionospheric Faraday rotation by more than 0.5 rad m$^{-2}$. 
Horologium field

<table>
<thead>
<tr>
<th>Central position</th>
<th>$(l, b) = (137^\circ, 7^\circ)$</th>
</tr>
</thead>
<tbody>
<tr>
<td>Size</td>
<td>$\sim 7^\circ \times 7^\circ$</td>
</tr>
<tr>
<td>Pointings</td>
<td>$5 \times 5$</td>
</tr>
<tr>
<td>Frequencies</td>
<td>$341, 349, 355, 360, 375 \text{ MHz}$</td>
</tr>
<tr>
<td>Resolution</td>
<td>$5.0^\prime \times 5.0^\prime \sec \delta = 5.0^\prime \times 5.5^\prime$</td>
</tr>
<tr>
<td>Noise</td>
<td>$\sim 5 \text{ mJy/beam (0.7 K)}$</td>
</tr>
<tr>
<td>Conversion Jy-K</td>
<td>$1 \text{ mJy/beam} = 0.146 \text{ K (at 350 MHz)}$</td>
</tr>
</tbody>
</table>

<table>
<thead>
<tr>
<th>Spacings</th>
<th>36m</th>
<th>48m</th>
<th>60m</th>
<th>72m</th>
<th>84m</th>
<th>96m</th>
</tr>
</thead>
<tbody>
<tr>
<td>Observing date</td>
<td>95/12/19</td>
<td>95/12/20</td>
<td>95/12/27</td>
<td>96/01/02</td>
<td>95/12/12</td>
<td>96/01/08</td>
</tr>
<tr>
<td>Start time (UT)</td>
<td>15:12</td>
<td>14:53</td>
<td>14:20</td>
<td>14:02</td>
<td>16:42</td>
<td>13:38</td>
</tr>
<tr>
<td>End time (UT)</td>
<td>02:58</td>
<td>02:53</td>
<td>02:20</td>
<td>02:02</td>
<td>03:25</td>
<td>01:38</td>
</tr>
</tbody>
</table>

Table 5.1: Observational details of the Horologium field

5.2.1 Missing large scale structure

An interferometer is insensitive to structure on large angular scales due to missing short spacings. The smallest baselines attainable with the WSRT of 36m means that scales above approximately a degree are attenuated sufficiently as to be undetectable. The $Q$ and $U$ maps are constructed so that in each mosaic pointing, the map integrals of $Q$ and $U$ are zero. This leads to missing large-scale components in $Q$ and $U$, and therefore erroneous determinations of $P$, $\phi$ and $RM$. However, if the variation in $RM$ is large enough within one pointing, the variation in polarization angle is so large that the average $Q$ and $U$ are close to zero. In this case, missing large-scale components are negligible (Chapter 3). In the field of observation discussed here, $\sigma_{RM} > 1.8 \text{ rad m}^{-2}$ in most pointings, which means that in these pointings missing large-scale components cannot account to more than a few percent of the small-scale signal. In four pointings, $\sigma_{RM} \approx 1.3 - 1.6 \text{ rad m}^{-2}$, which means that offsets could amount to 15 - 30% of the signal. Therefore, care must be exercised in interpreting the polarization data from one of the pointing centers which show $\sigma_{RM} \leq 1.8 \text{ rad m}^{-2}$. On the other hand, polarization angles show a linear variation over frequency in a large part of the data, yielding good $RM$ determinations, which would not be possible if large-scale offsets dominate. See Chapter 3 for an extended discussion of this point.

5.3 Observational results

5.3.1 Stokes $Q$ and $U$

The observed Stokes $Q$ and $U$ intensities at the $3^\prime$ resolution are shown in Fig. 5.1. The distributions of $Q$ and $U$ are approximately Gaussian, centered around zero, and have a width of 20 to 25 mJy/beam (equivalent to polarized brightness temperatures $T_{b,\text{pol}}$ of 2.9 - 3.7 K) for the five frequencies. The ring structure is clearly visible in $Q$
and $U$. To emphasize the perfect circularity of the ring, a circle of radius 1.44° and centered on $(\alpha, \delta) = (48.05°, 65.73°)$ is superimposed.

5.3.2 Polarized intensity $P$

The maps of polarized intensity $P$ derived from the Stokes $Q$ and $U$ maps are shown in Fig. 5.2 for all 5 frequencies. White denotes the highest intensity, and intensities above 110 mJy/beam (equal to $T_{\text{B, pol}} = 16$ K) are saturated. The maximum intensities in the 5 frequency bands are 141 mJy/beam, 122 mJy/beam, 107 mJy/beam, 120 mJy/beam, and 143 mJy/beam, respectively. Superimposed are lines of constant Galactic latitude $b = 3°$, 5°, 7°, 9°, and 11°, and the superimposed circle is the same as the one in Fig. 5.1. Several regions of different topology of polarized intensity are present in the field:

1. The most conspicuous feature is the ring in the center of the field. Although it is not as distinct as in polarization angle and Stokes $Q$ and $U$, the ring is still clearly visible in $P$ over most of its circumference. Only in the southwest, the ring is not well defined. At 375 MHz, the ring seems to be more blurred and unsharp than at the other frequencies.

2. A linear structure of high polarized intensity extends from the center of the ring towards the northwest, and is approximately aligned with Galactic latitude. We discuss this elongated structure of high $P$ in Section 5.7.1.

3. The southwestern corner of the field shows a filamentary pattern of “canals” of low $P$ running from southeast to northwest, approximately along lines of constant
Figure 5.2: Polarized intensity $P$ for each frequency, at $\delta$ resolution, and total intensity $I$ in the lower right plot. Superimposed are lines of constant Galactic latitude, and the same circle as in Fig. 5.1.
Galactic latitude, that are remarkably straight over many degrees. This pattern is also visible in angle (Fig. 5.3) and Stokes $Q$ and $U$ (Fig. 5.1). The southwestern part of the ring appears deformed in the direction of the filaments, but this deformation can be caused by averaging over the line of sight, so that the ring and filaments are not necessarily located at the same distance. Across the canals, there is an angle difference of $90^\circ$ (or $270^\circ$, $450^\circ$ etc.), indicating that these canals are caused by beam depolarization at the boundary of two regions of constant angle (Haverkorn et al. 2000, Chapter 3 and Chapter 6). These long canals are at the same location at all frequencies, but do not always have the same depth.

4. In the north, a patchy structure of polarization with an angular scale of $\sim 1^\circ$ is visible, spreading over the northern half of the field of observation.

5. In the southeast, a large region of very low polarized intensity exists, which could be due either to a region of low polarized emission or a region of large depolarization.

5.3.3 Total intensity $I$

In the lower right plot of Fig. 5.2 a map of total intensity $I$ at 349 MHz is shown, for which point sources $> 5$ mJy/beam are removed. The map has the same resolution of $5.0' \times 5.3'$ and the same brightness scaling as the $P$ maps in the other panels of the same figure.

No structure in total intensity $I$ is visible, although there is abundant structure in polarization. The circular structure in the upper left corner of the $I$ map is artificial and caused by a very bright unpolarized extragalactic source at that position. Due to the high-pass filter action of an interferometer, the map integral $I$ over the field is set to zero by lack of information about its true level. From the single dish survey of Haslam et al. (1981, 1982) at 408 MHz, the total brightness temperature at 408 MHz at this position is approximately 44.5 K with a temperature uncertainty of $\sim 10\%$ and including 2.7 K from the cosmic microwave background. To obtain the brightness temperature of the diffuse emission at 350 MHz, the contribution of 2.7 K from the CMBR is subtracted, as well as the contribution of 25% from discrete sources, derived from source counts (Bridle et al. 1972), assuming the spectral index of the Galactic background and the extragalactic sources to be identical. The intensities from the Haslam survey were scaled to our frequencies with a temperature spectral index of $-2.7$. We thus estimate the total brightness temperature at $\sim 350$ MHz to be 47 K. The apparent degree of polarization $p$, i.e. the polarized intensity divided by the total intensity, is mostly above 100%. This indicates that $I$ is uniform on the scales that the WSRT is sensitive to, and therefore the measured $I$ is close to zero. Small-scale structure in polarization due to Faraday rotation and depolarization causes higher $P$ than $I$, resulting in $p > 100\%$.

5.3.4 Polarization angle

In Fig. 5.3 we show a grey scale plot of polarization angle in the Horologium region at 349 MHz. The range in angle is $[-90^\circ, 90^\circ]$, so that white denotes the same angle
as black. The ring-like structure is well visible, and is close to circular everywhere except on the south-western side. The angle coherence extends to a diameter of \(\sim 4^\circ\). The linear structures in angle in the south-western corner of the field are aligned with Galactic latitude (see also Fig. 5.2).

The variation of polarization angle across the ring is shown in Fig. 5.4, where the grey scale is \(P\) at a frequency of 349 MHz and the polarization (pseudo-)vectors are superimposed. The length of the vectors is proportional to \(P\). The superimposed contours are contours of polarization angle, and the circle is the same circle as in Fig. 5.1. Note the remarkably constant polarization angle over the surface of the ring, apparent in both figures. The only place where the angle is disturbed is in the south-west.

### 5.3.5 Rotation measure

Rotation measures were calculated by linearly fitting the observed polarization angle \(\phi\) as a function of \(\lambda^2\). The polarization vectors have a \(\pm 180^\circ\) ambiguity, but the derived rotation measures are so small that this ambiguity does not play a role in our observations: \(|RM| \leq 10\ \text{rad m}^{-2}\), which means an angle change over the full frequency range of 341 MHz to 375 MHz \(\Delta \phi \leq 75^\circ\). Therefore, we calculate the \(RM\) values straightforwardly with angle differences minimized.

There are various mechanisms that can destroy the linear relation between \(\phi\) and \(\lambda^2\) and make \(RM\) determination unreliable, such as depolarization mechanisms (see Section 5.5) and the insensitivity to large-scale structure (see Section 5.2.1). We have only taken into account those \(RM\) values for which the linear \(\phi(\lambda^2)\)-relation has \(\chi^2_{\text{red}} <\)
Figure 5.4: Detail of the Horologium field, where superimposed vectors are polarization vectors, and the grey scale is polarized intensity at 349 MHz. The superimposed circle is the same as in Fig. 5.1. The contours delineate also polarization angle, showing that polarization angle is constant over the ring, except for the perturbation in the southwestern corner.

2 and where the polarized intensity averaged over all wavelengths $\langle P \rangle > 20$ mJy/beam ($\sim 4 \sigma$), which is $\sim 31.3\%$ of the data. However, of all pixels with $\langle P \rangle > 20$ mJy/beam, $\sim 62\%$ has a $\chi^2_{red} < 2$. The resulting $RM$ map is given in Fig. 5.5. The circles denote valid $RM$ values according to the above definition, where the diameter of the circle is proportional to the magnitude of the $RM$. Filled circles denote positive rotation measures. For clarity, only one out of two independent beams in both directions is shown.

In the area enclosed by the ring all the $RM$s are negative. The average $RM$ at the eastern part of the ring is about $-3$ rad m$^{-2}$, and decreases towards the center of the ring to $RM \approx -8$ rad m$^{-2}$. The patch of high $P$ at $(\alpha, \delta) = (51.5^\circ, 64^\circ)$ shows an approximately constant $RM$ of $-1$ rad m$^{-2}$, and further away from the ring, $RM$s increase up to a few rad m$^{-2}$ at the edges of the field.

These values are in agreement with earlier measurements at lower resolution. In rotation measure maps produced by Bingham and Shakeshaft (1967) and by Spoelstra (1984), $|RM| \leq 3$ rad m$^{-2}$ outside the ring. They presented values of $RM$ inside the ring of $RM < -5$ rad m$^{-2}$ and $RM < -3$ rad m$^{-2}$, respectively.
Figure 5.5: Rotation measure map. Circles represent RMs for which reduced \(\chi^2 < 2\) and \(\langle P \rangle > 20 \text{ mJy/beam} \approx 4\sigma\). The diameters of the circles scale with RM, and positive RMs are denoted by filled circles. Only one in four beams is shown.

5.3.6 Extragalactic sources

Twelve polarized extragalactic sources were detected in the field. The properties of the extragalactic sources were measured in maps having a resolution of 1\('\). All sources have a polarized intensity greater than 4 mJy/beam in each frequency band (the noise of the 1\('\) resolution data is about 1 mJy/beam) and a degree of polarization greater than 1%. Near the outer edge of the outermost pointing centers, instrumental polarization increases considerably, so that sources in that region were excluded. The characteristics of the polarized extragalactic sources are shown in Table 5.2, the \(\phi(\lambda^2)\)-fits in Fig. 5.6, and the positions of the sources in Fig. 5.7.

The brightest polarized source in the field with high polarization is the giant double lobed radio galaxy WNB 0313+683 at \((\alpha, \delta) = (48.25^\circ, 68.3^\circ)\) (Schoenmakers et al. 1998). In our analysis this extended source was detected in four separate maxima, viz. sources 5, 6, 7, and 8, all located at approximately the same position in Fig. 5.7. The high resolution measurements of Schoenmakers et al. (15\('\) in the RM map) show an average \(RM = -10.64 \text{ rad m}^{-2}\), which they argue is Galactic, and a residual \(RM\) of about 2 rad m\(^{-2}\) in each of the lobes. In our 1\('\) resolution observations at 350 MHz, this would result in a variation in polarization angle within the beam. Depolarization due to this variation in polarization angle can destroy the linear relation between \(\phi\)
and $\chi^2$ and result in the relatively high $\chi^2$ values in sources 5 and 8.

Source no. 3 is the only source with a positive $RM$, which is derived from a good fit with $\chi^2_{red} = 0.84$. As a check, taking into account the $180^\circ$ ambiguity could not yield a $RM$ value closer to that of the other sources. The next best fit has $\chi^2_{red} = 220$, so $RM = 1.6 \text{ rad m}^{-2}$ is indeed the only acceptable $RM$ determination for this source.

### 5.4 Observed properties of the ring-like structure

The northeastern half of the ring is very regular in polarized intensity, polarization angle and $RM$, whereas the southwestern part appears to be influenced by the linear structure aligned with Galactic latitude. Therefore, in analyzing the ring structure, we study radial averages of $P$, $\phi$ and $RM$ over position angles (N through E) $-20^\circ < \theta < 135^\circ$, i.e. the undisturbed part of the ring. The center of the radial averages is at $(\alpha, \delta) = (48.65^\circ, 65.73^\circ)$, which is the center of the ring superimposed in Fig. 5.1.

In Fig. 5.8, we show the radial averages of observed quantities out to a radius of about twice as large as the ring. The top panels show $P$ and $I$ at 5 frequencies. The peak in $P$ at a radius of about $1.4^\circ$ coincides with the position of the ring. The polarized brightness temperature at that radius is approximately 8 K, $T_{b,\text{pol}}$ inside the ring is about 3.8 K, and outside the ring about 3 K. The total intensity $I$ is completely uncorrelated with $P$, and hardly exceeds the noise. Note that the position and width of the peak in $P$ is frequency-independent, while inside the ring, the $P$ distribution varies with frequency.

The bottom left panel in Fig. 5.8 shows the polarization angle plotted as a function of radius, again at 5 frequencies. The angle gradient is remarkably linear between a radius of $0.6^\circ$ and $1.7^\circ$, with some structure at a radius of $\sim 1^\circ$. Whereas $P$ has a maximum at $\sim 1.4^\circ$, the linear increase of polarization angle with radius continues to
Figure 5.7: Rotation measures of polarized extragalactic sources indicated by white circles are overlaid on a grey scale representation of polarized intensity at 340 MHz. The $RM$ scales with the diameter of the circles, where the extrema are $-52.6$ rad m$^{-2}$ and $1.6$ rad m$^{-2}$, and open circles denote negative $RM$. Instrumental polarization increases rapidly beyond the outermost pointing centers, denoted by crosses.

Table 5.2: Polarization data for extragalactic sources with measured polarization in the Horologium field. The second column gives position, and the third $RMs$ with errors. Reduced $\chi^2$ of the $\phi(\lambda^2)$-relation is given in column 4. Columns 5, 6 and 7 give resp. $P$, $I$ (both in mJy/beam) and degree of polarization $p$ in percents, averaged over the five frequency bands.

<table>
<thead>
<tr>
<th>No.</th>
<th>($\alpha, \delta$) (°, °)</th>
<th>$RM$ (rad m$^{-2}$)</th>
<th>$\chi^2$</th>
<th>$\langle P \rangle$</th>
<th>$\langle I \rangle$</th>
<th>$\langle p \rangle$</th>
</tr>
</thead>
<tbody>
<tr>
<td>1</td>
<td>[52.83, 65.20]</td>
<td>$-16.2 \pm 1.2$</td>
<td>0.37</td>
<td>4.8</td>
<td>168</td>
<td>2.9</td>
</tr>
<tr>
<td>2</td>
<td>[51.90, 64.64]</td>
<td>$-17.8 \pm 1.2$</td>
<td>0.67</td>
<td>4.5</td>
<td>253</td>
<td>1.8</td>
</tr>
<tr>
<td>3</td>
<td>[50.72, 65.53]</td>
<td>$1.6 \pm 0.6$</td>
<td>0.84</td>
<td>8.7</td>
<td>97</td>
<td>9.0</td>
</tr>
<tr>
<td>4</td>
<td>[48.99, 66.21]</td>
<td>$-37.4 \pm 1.2$</td>
<td>1.67</td>
<td>4.4</td>
<td>73</td>
<td>6.0</td>
</tr>
<tr>
<td>5</td>
<td>[48.35, 63.33]</td>
<td>$-11.5 \pm 0.3$</td>
<td>3.55</td>
<td>15.5</td>
<td>84</td>
<td>19.2</td>
</tr>
<tr>
<td>6</td>
<td>[48.33, 68.28]</td>
<td>$-9.1 \pm 0.5$</td>
<td>0.47</td>
<td>10.4</td>
<td>88</td>
<td>12.1</td>
</tr>
<tr>
<td>7</td>
<td>[48.20, 68.27]</td>
<td>$-13.3 \pm 0.4$</td>
<td>1.61</td>
<td>10.6</td>
<td>72</td>
<td>15.0</td>
</tr>
<tr>
<td>8</td>
<td>[48.24, 68.25]</td>
<td>$-10.8 \pm 0.1$</td>
<td>13.34</td>
<td>37.3</td>
<td>366</td>
<td>10.1</td>
</tr>
<tr>
<td>9</td>
<td>[46.16, 63.88]</td>
<td>$-52.6 \pm 1.2$</td>
<td>0.65</td>
<td>4.4</td>
<td>82</td>
<td>5.3</td>
</tr>
<tr>
<td>10</td>
<td>[45.49, 67.54]</td>
<td>$-43.5 \pm 0.9$</td>
<td>1.05</td>
<td>6.4</td>
<td>218</td>
<td>2.9</td>
</tr>
<tr>
<td>11</td>
<td>[44.15, 64.24]</td>
<td>$-51.1 \pm 0.8$</td>
<td>1.53</td>
<td>5.6</td>
<td>364</td>
<td>1.0</td>
</tr>
<tr>
<td>12</td>
<td>[43.85, 68.48]</td>
<td>$-12.8 \pm 1.1$</td>
<td>1.56</td>
<td>5.5</td>
<td>238</td>
<td>2.3</td>
</tr>
</tbody>
</table>
at least $1.7^\circ$. This indicates that the structure or feature responsible for the ring in $P$ may be actually larger than the apparent size of the ring.

The bottom right panel gives the radially averaged observed $RM$ (only values with $\chi^2_{red} < 2$ and $(P) > 20$ mJy/beam). The errors in the $RM$ are the errors in the mean of the distribution of $RM$ over the indicated radial interval. The $RM$ does not vary much over the region at which the peak in $P$ occurs. At radii larger than $1.7^\circ$, the $RM$ is slightly positive: $0 \leq RM \leq 2$ rad m$^{-2}$. At smaller radii, $RM$ is negative, decreasing to $RM \approx -8$ rad m$^{-2}$ in the center, with a small increase around a radius $r = 0.7^\circ$.

The observations in $I$, $P$, $\phi$ and $RM$ impose stringent constraints on the nature of the ring. The distribution of $RM$ provides the strongest constraint: $RM$ is positive outside the ring, negative inside and the most negative at the center. Although $RM$ structure is created by a combination of structure in thermal electron density and magnetic field, a change in sign of $RM$ always indicates a reversal of the parallel component of the magnetic field. So the electron density and/or magnetic field configuration that causes the ring must reverse magnetic field directions from outside to inside the ring, and the magnetic field and/or density has to be the highest at the center. Furthermore, the lack of correlated $I$-structure implies that the ring in $P$ cannot have been produced by emission. In the next section, we discuss what processes can cause the observed distributions of $P$, $\phi$ and $RM$ in the ring. Subsequently, we shall describe in Section 5.6 some known structures and objects in the ISM, and discuss whether these are related to the ring structure.

5.5 The nature of the ring in $P$, $\phi$ and $RM$

From comparison of the $I$ and $P$ maps in Fig. 5.8, it is clear that the structure in $P$ cannot, even in part, be caused by structure in $I$. Structure in $P$ can also be created by missing large-scale structure in $Q$ and/or $U$ (see Chapter 3), but in Section 5.2.1 we have shown that in these observations missing large-scale structure cannot dominate. Therefore, the ring in $P$ must be due to a lack of depolarization. Several depolarization mechanisms can contribute to create the ring in $P$. We shall discuss briefly the different depolarization mechanisms thought to be of importance (for details see Chapter 3).

5.5.1 Depth depolarization

Depth depolarization is defined as all depolarization processes occurring along the line of sight and can be due to different physical processes. First, if the magnetic field in a synchrotron-emitting medium has small-scale structure, then the emitted (intrinsically) polarization angle of the synchrotron radiation will vary along the line of sight, causing wavelength independent depolarization. Secondly, if the medium also contains thermal electrons, the polarization angle of the radiation will be modulated by Faraday rotation as well, which causes additional depolarization (internal Faraday dispersion). So small-scale structure in (parallel) magnetic field and/or thermal electron density within the synchrotron emitting medium causes small-scale depolarization. These processes were discussed analytically by Sokoloff et al. (1998) for several different geometries of the medium, and numerically in Chapter 3 using observational constraints.
Figure 5.8: Radial averages of observed polarized intensity $P$ in Jy/beam for five frequencies (top left), total intensity $I$ in Jy/beam for five frequencies (top right), polarization angle $\phi$ in degrees for five frequencies (bottom left) and $RM$ (bottom right), averaged over the northern and eastern part of the ring structure (position angle $-20^\circ < \theta < 135^\circ$).

Figure 5.9: Radially averaged $P$ (left), $\phi$ (center) and $RM$ (right) plotted as a function of radius for a simple model of a $B$ and $n_e$ distribution. For details see text. The different lines refer to the 5 frequencies.
We modeled the effect of depth depolarization in the observations of the ringstructure using simple distributions of electron density $n_e$ and magnetic field $B$ on a rectangular grid. These distributions are not self-consistent, but the only goal of this simple model is to obtain a $P$ and $\phi$ distribution that is similar to the observations, i.e. approximately linear in $\phi$ and ring-like in $P$. Synchrotron radiation of emissivity $\varepsilon \propto B^2$ is emitted in the regions where $B_\perp$ is non-zero, and is Faraday-rotated while propagating through the medium, depending on the local $B_\parallel$ and $n_e$ distributions. Furthermore, a polarized background contribution $P_b$ is added, which is also Faraday-rotated. Both magnetic field components, parallel and perpendicular to the line of sight, were assumed constant within a circle of radius $r_0$, and to decrease to zero outside the circle as $r^{-5}$. The electron density distribution is highest in the center of the circle, and decreases as $r^{-0.4}$ outwards. These configurations were chosen because they approximately reproduce the observed $\phi$ and $RM$ distributions, as is shown in Fig. 5.9.

This figure shows the model output $P$ and $\phi$ at 5 frequencies, and $RM$. We have chosen $B_\parallel = -3.5$ $\mu$G, $B_\perp = -2$ $\mu$G, $n_e = 0.2$ cm$^{-3}$, and $P_b = 4$ K. This reproduces the shape and magnitude of $\phi$ and $RM$ reasonably well, but the $P$ distribution is very different from the observed one. First, the predicted $P$ at the center is much larger than is observed. However, this discrepancy could be explained by assuming a chaotic magnetic field component at the center of the circle (without worrying yet what this could mean physically). This would result in depolarization and a lower observed $P$ in the center of the modeled circle.

However, a more severe problem is posed by the wavelength dependence of the model predictions. Although our models have very different $B$ and $n_e$ distributions and either spherical or cylindrical symmetry, they all show a distinct wavelength dependence of the peak in $P$, as in Fig. 5.9. But from the observations, the position of the peak in $P$ does not change with wavelength, see Fig. 5.8. The wavelength dependence of $P$ appears to be a generic property of all models involving depolarization due to depth depolarization. However, the mechanism that can create wavelength independent depolarization, viz. tangled magnetic fields, yields structure in $I$, contrary to what is observed. Therefore we conclude that depth depolarization cannot be the main process that creates the ring in $P$, although we do expect depth depolarization to be present, e.g. in depolarizing the (uniform) background.

5.5.2 Beam depolarization

Beam depolarization, i.e. the averaging out of polarization vectors across one synthesized beam, is significant in the field. As there is structure in $RM$ on beam scales, it is likely that $RM$ varies on scales smaller than the beam as well. Furthermore, at the positions of the depolarization canals, the influence of beam depolarization is clearly visible. (Partial) beam depolarization can destroy the linear $\phi(\lambda^2)$-relation, but does not necessarily do so. At low polarized intensities, the influence of beam depolarization can be considerable, and observed $RM$ values at low polarized intensity should be used with care, as they can deviate from the true $RM$ value.

Beam depolarization, due to chaotic structure in polarization angle on scales smaller than the beam, can arise due to tangled magnetic fields and/or small-scale variations
in thermal electron density. A possible explanation for the lack of $P$ in the central part of the ring could be a chaotic magnetic field in the center, while the outer parts of the ring exhibit very coherent magnetic fields and electron density. However, although $P$ drops towards the center of the ring, polarization angles are coherent over the whole structure, indicating that beam depolarization due to chaotic sub-beam-scale structure does not dominate.

Beam depolarization can also be due to a spatial gradient in $RM$. For a gradient $dRM/dr$ over many beams, the depolarization factor $p$ is (Gaensler et al. 2001)

$$p = \frac{P_{\text{observed}}}{P_{\text{original}}} = \exp \left[-\frac{1}{\ln 2} \left(\frac{dRM}{dr}\right)^2 \lambda^4\right]$$

where $dRM/dr$ is the gradient of $RM$ over the beam and $r$ is the radius of the ring, expressed in the number of beams. In the upper panel of Fig. 5.10, we compare the observed radial gradient in $RM$ (solid line) with the theoretical $RM$ gradient from Eq. (5.1) (dotted line), using the observed polarized intensity $P$ at 341 MHz. Using $P$ from other frequency bands gives very similar results. The derivative of the observed $RM$ was computed after smoothing $RM$ by about 1.5 beams. At the position of the ring, the modeled $RM$ gradient shows the same decrease as the observed gradient. Furthermore, at the position where the gradient in $RM$ increases, the depolarization also increases. This indicates that the high $P$ within the ring, as well as the decrease in $P$ at the inner and outer boundaries, can be caused by the gradient in $RM$. However, in the center of the ring, as well as outside the ring, the existence of a gradient in $RM$ would result in a lower depolarization than is observed. So both at the center and outside the ring, other depolarization mechanisms must be important too.

5.6 Connection of the ring with known ISM objects

The distance to the ring-like structure is not constrained. Depth and beam depolarization introduce a Faraday depth or depolarization horizon, because polarized radiation emitted at large distances is depolarized more than radiation emitted nearby (Chapter 3). But because the foreground and background of the ring must be very uniform, the polarization horizon can be at very large distances. However, if the ring is located at large distances, say in or behind the Perseus arm, it would require a path length of more than 2 kpc with an unusually uniform magnetic field and electron density along the path length. Moreover, if the ring is in the Perseus arm, its size would be $>100$ pc, which is unlikely in view of the regular shape of the ring. Therefore, the ring is unlikely to be located in (or behind) the Perseus arm and is probably an inter-arm feature.

The reversal in the direction of the magnetic field from outside to inside the ring, and the $RM$ that becomes more negative towards the center, constrain possible explanations for the ring strongly, because they would not allow magnetic field configurations that have spherical symmetry (such as stellar winds or supernovae). This is because in those configurations, the $RM$ contribution from the front half of the structure compensates the (opposite) $RM$ contribution from the far end (e.g. radial fields in young supernovae, bubbles blown in a regular magnetic field perpendicular to the line of sight). Alternatively, the $RM$ would be higher at the edges of the ring than in the center (e.g.
bubbles blown in a regular magnetic field parallel to the line of sight). Instead, the structure that creates the ring-like structure, must have a magnetic field directed away from us, in surroundings where the magnetic field has a small component towards us. Furthermore, the magnetic field strength and/or electron density must increase from the edge towards the center of the ring. These constraints make a connection of the ring with many known classes of objects unlikely. Below we discuss the plausibility that some known gaseous structures in the ISM could be responsible for the polarized ring.

**Planetary nebula** The circular form of the ring structure suggests it might be due to a planetary nebula (PN). The radio emission from a PN is free-free emission, which is negligible at our low frequencies for any reasonable temperature. However, the ring has an angular diameter more than ten times as big as any other PN observed before. If the PN were 5 pc in size, about the size of the largest ancient PNe known (Tweedy and Kwiter 1996), then it would be at a distance of about 100 pc and it is unlikely that it hasn’t been observed before. Furthermore, it is difficult to see how the magnetic field configuration needed to create the ring could be present in a PN. For these reasons we believe it unlikely that the ring structure is a planetary nebula.
Strömgren sphere from HD20336 Close to the projected center of the ring the B2V star HD20336 is located at a distance of 246 ± 37 pc (Hipparcos Catalogue, Perryman et al. 1997). Verschuur (1968) suggested that the ring could be related to this star. If the ring-structure in $P$ coincides with the Strömgren sphere of HD20336, its radius would be $R_s = 6$ pc. The excitation parameter $U = R_s n_e^{4/3}$ is 2.6 pc cm$^{-3}$ for a B2V star (Panagia 1973), which yields an electron density inside the Strömgren sphere of $n_e \approx 0.1$ cm$^{-3}$. The emission measure $EM = \int n_e^2 \, dl = 2U^3 / R_s^5$ solely from the Strömgren sphere is 0.96 cm$^{-6}$ pc ($\sim 2$ R), which would probably be detectable in the WHAM H$\alpha$ survey (see Section 5.7.3). In addition, the Strömgren sphere explanation is unlikely because of the high proper motion of the star, viz. 18 mas/yr, which is equivalent to 5$^5$ in a Myr. The recombination time $t_r \approx (\alpha(2)n_e)^{-1}$, where the recombination coefficient to the second level of the hydrogen atom $\alpha(2) \approx 1.4 \times 10^{-13}$. So the recombination time is 8 Myr with $n_e = 0.1$ cm$^{-3}$. Due to the high proper motion of the star, we conclude that a circular Strömgren sphere cannot be maintained. Rather, the structure would be elongated in the opposite direction to the proper motion of the star. In addition, it is difficult to imagine a Strömgren sphere with a sufficiently asymmetric magnetic field structure to explain the observed $RM$, as described above. Hence, the ring is unlikely to be a Strömgren sphere around the star HD20336.

Stellar wind from HD20336 We use the interstellar wind-blown bubble model by Weaver et al. (1977) to estimate the time needed to blow a bubble of radius 6 pc with a stellar wind from a B2V star. Weaver et al. estimate the radius of a blown bubble as

$$R_s = 26.5 \left( L_{36} n_H^{-1} t_6^{1.3} \right)^{0.2} \text{ pc}$$

where $L_{36}$ is the mechanical energy of the stellar wind in units of $10^{36}$ erg s$^{-1}$, $n_H$ is the original hydrogen particle density before passage of the shell, and $t_6$ is the elapsed time in Myrs. Assuming that $L_{mech} = 2.5 \times 10^{34} L_{bol}$ (Israel and van Driel 1990), and that the absolute luminosity of a B2V star is $L = 3.7 L_\odot$ (Panagia 1973), then $L_{36} = 2.2 \times 10^{-4}$ erg s$^{-1}$. With $R_s = 6$ pc and $n_H = 0.1$ cm$^{-3}$, the time needed to blow a bubble of the size of the ring is almost 5 Myr. If $n_H$ is ten times lower, $t_6$ would still be 2.2 Myr. Compared to the proper motion of the star, the time needed to blow a circular bubble is far too long to produce the observed structure. Therefore, the ring cannot be due to the effect of a stellar wind either.

Supernova remnant For supernova remnants, a $\Sigma$-D relation exists between brightness and distance. Using the non-detection of signal in $I$ as an upper limit for the emission of the ring structure, the $\Sigma$-D relation of Case and Bhattacharya (1998) would imply a distance larger than 36 kpc if the ring was a SNR. Clearly the ring structure is much too faint in $I$ to be related to a supernova remnant.

Superbubble or chimney Superbubbles that become chimneys as they blow out material from the spiral arm where they originate, have been observed in HI, not only into the Galactic halo, but also into the inter-arm region in the Galactic plane (McClure-Griffiths et al. 2002). The ring could be such a chimney blown away straight from the Sun. Galactic magnetic field frozen in in the plasma that is blown out can
Table 5.3: Comparison of the ring discussed here with two earlier detections of circular or elliptical polarization structures by Gray et al. (1998) and Uyaniker and Landecker (2002).

<table>
<thead>
<tr>
<th>$l, b$</th>
<th>$91.8\deg, -2.5\deg$</th>
<th>$137\deg^2, 7\deg^2$</th>
</tr>
</thead>
<tbody>
<tr>
<td>size</td>
<td>$10^\circ \times 2^\circ$</td>
<td>$2.8^\circ \times 2.8^\circ$</td>
</tr>
<tr>
<td>$\Delta \phi$</td>
<td>280$^\circ$</td>
<td>300$^\circ$</td>
</tr>
<tr>
<td>$\Delta RM$</td>
<td>110 rad m$^{-2}$</td>
<td>40 rad m$^{-2}$</td>
</tr>
<tr>
<td>$\phi$ structure</td>
<td>linearly increasing</td>
<td>“linearly” decreasing</td>
</tr>
<tr>
<td>$P$ structure</td>
<td>ring</td>
<td>ring</td>
</tr>
<tr>
<td>distance</td>
<td>$440$ pc $&lt; d &lt; 1.5$ kpc</td>
<td>$350 \pm 50$ pc</td>
</tr>
</tbody>
</table>

cause the observed high negative $RM$. However, if this is the case, one would expect the magnetic field to be maximal at the edges of the chimney, and a low electron density inside the chimney.

**Magnetic structure**

It is possible that the ring is created through bending of magnetic field lines. E.g. contraction and motion of a plasma cloud or an analogue to the Parker instability can enhance and reverse magnetic fields. The $RM$ plot in Fig. 5.8 shows an increase in $RM$ out to a radius of about 3$^\circ$. If the radial average over $RM$ is taken over 360$^\circ$, $RM$ still increases slowly to 0 rad m$^{-2}$ at $r = 3^\circ$ and to 2 rad m$^{-2}$ at $r = 4^\circ$. This behavior is more like a magnetic structure, with decreasing magnetic field over a large radius, than like a structure in thermal electron density which has more or less definite boundaries. We conclude that the ring-like structure in $P$ must be a funnel-shaped enhancement of magnetic field, directed straight away from us, possibly related to magnetic flux tubes (Parker 1992, Hanasz and Lesch 1993).

A “magnetic anomaly” of a few degrees in size has been found in the direction $l \approx 92^\circ$, $b \approx 0^\circ$ (Clegg et al. 1992, Brown and Taylor 2001) from $RM$s of extragalactic sources. However, $RM$s of extragalactic sources in the direction of the ring do not show a local magnetic field reversal (Fig. 5.7). Therefore we expect the ring-structure to be more localized or less strong than this magnetic anomaly.

**5.6.1 Are there more ring-like structures?**

There have been two earlier detections of elliptical structures in polarization angle and polarized intensity without correlated structure in total intensity, both in single-frequency observations at 1420 MHz. The first detection by Gray et al. (1998) is an elliptical feature of about $2^\circ \times 1^\circ$ in size, which shows a linear increase in polarization angle towards the center. Polarized intensity shows a ring-like behavior, and the regular structure in $\phi$ extends to a larger radius than where $P$ peaks.

The “Polarization Lens” observed by Uyaniker and Landecker (2002) shows approximately similar characteristics: an approximately linear decrease of polarization angle towards the center, and a ring-like structure in $P$, although the regular structure in angle seems to trace an ellipse rather than a circle. The characteristics in $P$ and $\phi$ are so similar for the three features, that it is tempting to consider them as members of...
the same class of objects.

However, although the appearance of the three rings is similar, the changes in $RM$ responsible for the angle changes are vastly different. Although the first two objects are only detected at one frequency, and therefore their $RM$ cannot be determined, a spatial variation in polarization angle is equivalent to a change in $RM$. For the first detection $\Delta RM \approx 110 \text{ rad m}^{-2}$, $\Delta RM \approx 40 \text{ rad m}^{-2}$ for the second, and $\Delta RM \approx 8 \text{ rad m}^{-2}$ for the multi-frequency detection discussed here. If the objects are related, there are two explanations for the differences in $RM$. First, the range in $\Delta RM$ could be due to a difference in magnetic field strength and/or electron density, which would imply that the contraction of magnetic fields and/or enhancement in electron density creating the ring must vary over a huge range of scales. Secondly, the range in $\Delta RM$ could be due to a difference in size over which the RM exhibits a linear gradient with equal $\Delta RM$ per parsec for each ring. If we assume a distance for the ring detected by Gray et al. to be 1000 pc, then the first two objects would be consistent with a gradient of $\sim 3 \text{ rad m}^{-2}$ per parsec across the plane of the sky. If our observed ring would be a feature similar to this, the size of the ring would be about 3 pc and its distance 60 pc. In this scenario, rings with similar magnetic field and electron density characteristics would show a range in size (35 pc, 12 pc and 3 pc).

The first two objects were interpreted to be mainly due to an increase of electron density. This explanation cannot apply to the ring discussed here, however, because we observe a change in sign in $RM$ from outside to inside the ring, implying that the cause of the structure must be at least partly magnetic.

The reason that Gray et al. and Uyanker et al. assume that the rings reflect an increase in electron density are the following. First, if the structure was magnetic, the magnetic field enhancement $\Delta B_{||}$ would be $\sim 5$ times the random component of the magnetic field. This high magnetic field, and a configuration with the magnetic field directed along the line of sight was considered to be very unlikely.

However, in their estimate of the $\Delta B_{||}$ that is necessary to produce the observed $\Delta RM$, they assume that the ring-like structure is an oblate ellipse. If the magnetic field is funnel-shaped instead, the path length through the structure would increase enormously and lower $\Delta B_{||}$ values can account for the observed $\Delta RM$. An alignment with the line of sight seems fortuitous, but it may well be that it is the only configuration that we can observe, because only such a configuration has a long path length through a strong $B_{||}$, making $RMs$ dominant over the background.

We conclude that the magnetic field must play an essential role in the shaping of the rings, although an increase in electron density is likely to also be of importance. The three detected ring-like structures could be similar objects, although they would need to exhibit a range in magnetic field and electron density, and/or size.

### 5.7 Structure outside the ring

#### 5.7.1 The elongated structure of high $P$

The elongated structure of high $P$ extending from the center of the ring to the northwest agrees with the scenario sketched by Verschuur (1969). Using the Green Bank 300ft telescope, he observed a filamentary deficit of HI, where the filament starts at the
position of the B2V star discussed in Section 5.1, and extends in the direction opposite to the star’s proper motion, along the bar of high $P$ present in our observations. He argues that the star has tunneled through the HI and has blown the neutral material away. Radiation from the star could then have ionized the remaining low density material in the tunnel. This picture is consistent with the filament of high polarization that we see at the same position, from the center of the ring to the northwest.

However, the HI deficit seems to follow the ring of high $P$ as well at the western edge. If this is so, then the density inside the ring should be lower than the surrounding density, so that neutral HI is blown out of the ring more easily. In this case, the ionization tunnel and the ring structure are at the same distance (viz. $\sim 250$ pc), although they do not necessarily have the same origin.

### 5.7.2 Uniform Galactic magnetic field

The coherence in $\phi$, and to lesser extent in $P$, over a large part of the ring indicates that the polarized foreground or background have to be uniform on the scale of the ring. If structure in $RM$ on smaller scales was present somewhere in the line of sight, be it in front of or behind the ring-like structure, the linearity of angle with radius would be destroyed. Small deviations from linearity are present (see e.g. the patch of less negative $RMs$ at $(\alpha, \delta) = (49.5^{\circ}, 66^{\circ})$), which indicates a change in $RM$ of a few rad m$^{-2}$ somewhere along the line of sight. However, in general the medium must be very uniform to give such large coherent behavior in the polarization angle as in the ring.

The ring is located in the “fan region”, which is a region of high polarization even at low frequencies (Brouw and Spoelstra, 1976) and with very ordered polarization angles, located at about $120^{\circ} \leq \ell \leq 160^{\circ}$, $0^{\circ} \leq b \leq 20^{\circ}$. The high polarization in this region indicates less depolarization, and therefore the uniform magnetic field should dominate over the random magnetic field component. From external galaxies, there are indications that indeed the uniform magnetic field component is higher than the random component between the spiral arms (Beck and Hoernes 1996, Beck et al. 1996).

Polarized radiation originating at large distances from the observer will be depolarized due to its long path length through the depolarizing medium. In addition, beam depolarization becomes more severe at large distances because the angular scale of structure becomes smaller with distance. These two effects create an effective largest distance out to which polarized radiation can still be observed, the “polarization horizon”. This horizon is frequency dependent, and lies closer to the observer for lower frequencies. Because we observe at very low frequencies, we expect radiation from the Perseus arm to be completely depolarized, so that the polarization radiation observed comes only from the inter-arm region.

From Fig. 5.5, and the $RM$ determinations of Spoelstra (1984) and Bingham and Shakeshaft (1967), we derive that outside the ring, $RM \leq 6$ rad m$^{-2}$. Then, the parallel component of the magnetic field must be $B_\parallel \leq 0.2 \mu$G, assuming a path length of the polarized radiation of 600 pc (Chapter 3) and a thermal electron density of 0.08 cm$^{-3}$ (Reynolds 1991). The path length is probably longer than 600 pc, which only reduces $B_\parallel$ further. This very low value of $B_\parallel$ is in agreement with the uniform magnetic field being mostly perpendicular to the line of sight in this direction.
Figure 5.11: Left: $P$ at 349 MHz where the white box denotes the volume used in the right plot. Right: the symbols denote $RM$ against a coordinate in the approximate direction of Galactic longitude $l$ (top) and longitude $l'$ (bottom), where 1 pixel is $\sim 0.6\arcmin$. $RM$ is averaged over $l$ and $l'$, respectively. The plotted error in $RM$ is the error in the mean of $RM$ in each bin. The upper line is $P$ at 349 MHz averaged over the same bins.

5.7.3 The aligned depolarization canals

In the southwestern part of the region of observation, long and straight depolarization canals exist that are approximately aligned with Galactic latitude, see Fig. 5.2. Fig. 5.11 shows the distribution of $RM$ and $P$ along and across the depolarization canals, i.e. approximately along Galactic longitude $l$ and latitude $b$. The top plot shows the $RM$ distribution as a function of $b$, and the bottom plot as a function of $l'$, where the area used is indicated by the white box in the two left hand figures. $RM$ and $P$ are averaged in the direction perpendicular to $l'$ and $b$, respectively. The symbols in the right figures denote $RM$, where the error bars give the error in the mean of $RM$ in each bin. The upper lines denote the $P$ distribution at 349 MHz, averaged over a bin. The main difference between the two plots is that the $RM$ along the filaments hardly changes, whereas the $RM$ across the filaments shows a gradient and/or stratification.

To investigate whether this $RM$ structure is due to magnetic fields and/or thermal electron density, we can use independent measurements of $n_e$, e.g. from H$\alpha$ observations. We use data from the Wisconsin H$\alpha$ Mapper survey (WHAM, Haffner et al., in
prep., Reynolds et al. 1998), as shown in Fig. 5.12. The Hα intensity increases from \( \sim 2.2 \) R to 30 R (1 R = 1 Rayleigh is equal to a brightness of \( 10^6 / 4 \pi \) photons cm\(^{-2}\) s\(^{-1}\) ster\(^{-1}\), and corresponds to an emission measure \( EM \) of about 2 cm\(^{-6}\) pc for gas with a temperature \( T = 10000 \) K) in the velocity range \( v = -10 \) km s\(^{-1}\) to \( v = -50 \) km s\(^{-1}\), which coincides with distances of 300 pc to 2.3 kpc using the rotation curve of Fich et al. (1989). If the Hα increase is generated over a line of sight of 2 kpc, the necessary increase in thermal electron density is only from 0.055 cm\(^{-3}\) to 0.07 cm\(^{-3}\). If we take again the path length of the polarized radiation to be 600 pc, the change in \( RM \) seen in Fig. 5.11, viz 7 rad m\(^{-2}\), can be made with a constant \( B_0 \approx 0.2 \) \( \mu \)G, which agrees with the estimate made in Section 5.3.5 for the uniform magnetic field component.

But if the gradient in \( RM \) is constant, there would be no depolarization canals. So the \( RM \) increases discontinuously, as is visible in Fig. 5.11. If \( \Delta RM \) over one beam is between 2.1 and 2.5 rad m\(^{-2}\) (or between 6.3 and 7.5 rad m\(^{-2}\), etc.), depolarization canals occur. Fig. 5.13 shows \( P \) and \( RM \) over a narrow slit (\( \sim 1.5' \)) across the canals plotted against \( \phi \), located within the box in Fig. 5.11. The top panel shows plots for \( P \) at all 5 frequencies, oversampled by a factor of 5. The central panel shows \( RM \), where \( RM \) values are only plotted for beams for which \( \chi^2_{red} < 2 \) and \( \langle P \rangle > 20 \) mJy/beam. Polarization angle \( \phi \) is given in the bottom panel. The two sharpest depolarization canals visible at all frequencies, at pixel numbers 113 and 145 correspond to abrupt changes in \( RM \) of about 2 rad m\(^{-2}\) and angle changes around \( 90^\circ \). Abrupt \( RM \) changes with other magnitudes are present as well, but these do not create depolarization canals. For an extended discussion on the cause of the canals, see Chapter 3.
Figure 5.13: Spatial variations of $P$ at 5 frequencies (upper plot), $RM$ (central plot) and $\phi$ (lower plot) over a slit of $\sim 1.5'$ wide, against $\theta'$. $RM$ and $\phi$ values are only given where $\chi^2_{red} < 2$ and $\langle P \rangle > 20$ mJy/beam.

5.8 Conclusions

The ring-like structure in polarized intensity $P$ with a radius of about $1.4'$ shows a regular increase in polarization angle from its center out to $\sim 1.7'$, suggesting that the structure is circular instead of a ring, and extends to larger radii than the ring in $P$. The rotation measure is slightly positive outside the ring, reverses sign inside the ring and decreases almost continuously to $RM \approx -8$ rad m$^{-2}$ at the center. This property rules out its production by any spherically symmetric structure, such as a supernova remnant or a wind-blown bubble. The $RM$ structure, and the observation that the coherence in angle slowly disappears beyond a radius of $\sim 1.7'$ indicates a magnetic origin for the polarization ring, probably accompanied by an electron density enhancement. We propose that the ring is produced by a predominantly magnetic funnel-like structure, in which the magnetic field strength is maximal at the center and the parallel magnetic field is directed away from us. The enhancement in $P$ at radius $\sim 1.4'$ is caused by a lack of depolarization due to the relative constancy of $RM$ at that radius. A filamentary pattern of parallel, narrow depolarization canals indicates
structure in $RM$ which is aligned with Galactic longitude. The depolarization canals are probably created by abrupt spatial gradients in $RM$.

**Acknowledgements**

We are grateful to T. Spoeistra for allowing us to use his observations, and to R. Beck and E. Berkhuijsen for critical reading of the manuscript. The Westerbork Synthesis Radio Telescope is operated by the Netherlands Foundation for Research in Astronomy (ASTRON) with financial support from the Netherlands Organization for Scientific Research (NWO). The Wisconsin H-Alpha Mapper is funded by the National Science Foundation. MH is supported by NWO grant 614-21-006.

**References**

Beck R., & Hoernes P., 1996, Nat 379, 47
Berkhuijsen E. M., Brouw W. N., Muller C. A., & Tinbergen J., 1984, BAN 17, 465
Berkhuijsen E. M., & Brouw W. N., 1983, BAN 17, 185
Brulé A. H., Davis M. M., Fomalont E. B., & Lequeux J., 1972, NPhS 235, 123
Panagia N., 1973, AJ 78, 929
Chapter 5

Verschuur G. L., 1969, AJ 74, 597
Verschuur G. L., 1968, Obs 88, 15
Structure in the local Galactic ISM on scales down to 1 pc, from multi-band radio polarization observations


Abstract

We discuss observations of the linearly polarized component of the diffuse Galactic radio background. These observations, with an angular resolution of 4′, were done with the Westerbork Synthesis Radio Telescope (WSRT) in 5 frequency bands in the range 341 – 375 MHz. The linearly polarized intensity $P$ (with polarized brightness temperature going up to 10 K) shows a ‘cloudy’ structure, with characteristic scales of 15 – 30′, which contains relatively long, but very narrow ‘canals’ (essentially unresolved) in which $P$ is only a small fraction of that in the neighboring beams. These ‘canals’ are generally seen in more than one frequency band, although their appearance changes between bands. They are probably due to depolarization within the synthesized beam, because the change in polarization angle $\Delta \phi_{\text{pol}}$ across the deepest ‘canals’ is in general close to 90° (or 270° etc.). These very abrupt changes in $\phi_{\text{pol}}$, which are seen only across the ‘canals’, seem to be accompanied by abrupt changes in the rotation measure ($RM$), which may have the right magnitude to create the difference of close to 90° in $\phi_{\text{pol}}$, and thereby the ‘canals’. The structure in the polarization maps is most likely due to Faraday rotation modulation of the probably smooth polarized radiation emitted in the halo of our Galaxy by the fairly local ($\lesssim 500$ pc) ISM. Therefore, the abrupt changes of $RM$ across the ‘canals’ provide evidence for very thin ($\lesssim 1$ pc), and relatively long transition regions in the ISM, across which the $RM$ changes by as much as 100%. Such drastic $RM$ changes may well be due primarily to abrupt changes in the magnetic field.
6.1 Introduction

Wieringa et al. (1993) were the first to note structure on arcminute scales in the linearly polarized component of the Galactic radio background at 325 MHz, observed with the WSRT. The small-scale structure in the maps of polarized intensity \( P \) (with polarized brightness temperatures \( T_{b,\text{pol}} \) of up to 10 K), does not have a counterpart in total intensity, or Stokes \( I \), down to very low limits. Because the total Stokes \( I \) of the Galactic radio background has an estimated \( T_{b,\text{pol}} \) of the order of 30 – 50 K at 325 MHz, which must be very smooth and therefore filtered out completely in the WSRT measurements, the apparent polarization percentage of the small-scale features can become very much larger than 100%.

The absence of corresponding small-scale structure in Stokes \( I \) led Wieringa et al. (ibid.) to propose that the small-scale structure in polarized intensity \( P \) is due to Faraday rotation modulation. In this picture, synchrotron radiation generated in the Galactic halo reaches us through a magneto-ionic screen, viz. the warm relatively nearby ISM. Structure in the electron density and/or magnetic field in the ISM causes spatial variations in the rotation measure \( (RM) \) of the screen. Hence, the angle of linear polarization of the synchrotron emission from the halo is rotated by different amounts along different lines of sight. Even if the polarized emission in the halo were totally smooth, in intensity as well as angle, the screen would produce structure in Stokes \( Q \) and \( U \).

Small-scale structure in the polarized Galactic radio background has recently been observed also at other frequencies. At 1420 MHz, Gray et al. (1998, 1999) used the DRAO synthesis telescope to study the phenomenon at 1' resolution. Uyaniker et al. (1999) used the Effelsberg telescope at 1.4 GHz, to map the polarized emission at 9 resolution over about 1100'. Duncan et al. (1998) discuss radio polarization data at 1.4, 2.4 and 4.8 GHz with the Parkes radio telescope and the VLA, at resp. 3', 10' and 15' resolution. All these observations support the interpretation in terms of modulation of emission originating at larger distances, by a relatively nearby Faraday screen.

The distributions of polarized intensity and angle may therefore be used to study the structure of the Faraday screen. In particular, polarization observations give information about the electron density, \( n_e \), and the component of the magnetic field parallel to the line of sight, \( B_\parallel \), in the ISM on scales down to less than \( \sim 0.5 \text{ pc} \) (< 4') at an assumed distance of \( \sim 500 \text{ pc} \). The diffuse nature of the polarized radio background allows (almost) complete spatial mapping of \( RMs \) over large areas, provided one has observations at several frequencies. This gives a large advantage over \( RM \) determinations through individual objects, like pulsars or extragalactic radio sources.

6.2 Distribution of polarized intensity

In Fig. 6.1 we show a grey scale representation of the polarized intensity in a 5 MHz wide frequency band centered at 349 MHz. The map shows a region of \( 6.4' \times 9' \) centered at \( \alpha = 6h 10m, \delta = 33^\circ(\ell = 161^\circ, b = 16^\circ) \) at an angular resolution of about 4'. It is one of 8 frequency bands observed simultaneously. Three of those have strong interference, but we obtained good data at 341, 349, 355, 360 and 375 MHz. All 5 maps were made combining mosaics of 5×7 pointing centers. This yields constant sensitivity
Figure 6.1: Linearly polarized intensity $P$ at 349 MHz in a $6.4^\circ \times 9^\circ$ field centered at $\ell = 161^\circ, b = 16^\circ$. The resolution is $\sim 4'$, the maximum brightness temperature is $\sim 10$ K. The generally ‘cloudy’ distribution contains long narrow ‘canals’ of low $P$. The white box shows the area displayed in Fig. 6.2.

over a large area (see e.g. Rengelink et al. 1997). The observations were done with the WSRT in January and February 1996, largely at night, and ionospheric Faraday rotation was therefore well-behaved. No corrections were applied.

The region in Fig. 6.1 is rather special because $T_{b,\text{pol}}$ goes up to 10 K, and because it contains large, almost linear structures in $P$. Our attention was drawn to this field by the panoramic view of Galactic polarization produced in the WENSS survey (de Bruyn & Katgert 2000). However, this field is not unique, and there are other regions with similarly high $T_{b,\text{pol}}$. Over a very large fraction of the map the $P$-signal is quite
significant, with a noise $\sigma_T \approx 0.5$ K. With S/N-ratios of generally more than 3 and going up to 30, polarization angles are well-defined. Note that in this region, the upper limit to structure in Stokes $I$ (total intensity) on small scales ($\lesssim 30^\prime$) is about 1 K, or less than 2% of the total $I$.

There appear to be at least two distinct components in the polarized intensity distribution. The first one is a fairly smooth, ‘cloudy’ component, pervading the entire map, with intensity variations on typical scales of (several) tens of arcminutes. In addition, there are conspicuous, very narrow and often quite long and wiggly structures, which we will refer to as ‘canals’, in which the polarized intensity is considerably lower than in the intermediate surroundings. In this Letter we focus on the nature and implications of the narrow ‘canals’; we discuss the ‘cloudy’ component in more detail in a forthcoming paper (Haverkorn et al. 2003) and in Chapter 4.

### 6.3 The nature of the ‘canals’ in polarized intensity

The strong and abrupt decrease of polarized intensity in the ‘canals’ suggests that depolarization is responsible. There are several mechanisms that can produce depolarization, but the only plausible type in this case is beam depolarization. This occurs when the polarization angle varies significantly within a beam. Complete depolarization requires that for each line of sight there is a ‘companion’ line of sight within the same beam that has the same polarized intensity but for which the polarization angle differs by $90^\circ$. Below we will show that our observations indicate that the polarization angle indeed changes by large amounts across low polarized intensity ‘canals’, and close to $90^\circ$ across the ‘canals’ of lowest $P$.

Depolarization can also be caused by ‘differential Faraday rotation’. This happens when along a line of sight emitting and Faraday-rotating plasmas coexist (e.g. Burn 1966; Sokoloff et al. 1998). However, the absence of correlated structure in Stokes $I$ and the high degree of polarization suggest that this is not a dominating effect. Significant bandwidth depolarization, which occurs when the polarization angle is rotated by greatly different amounts in different parts of a frequency band could only play a role (given our 5 MHz bandwidth) if the $RM$ were of order $80 \text{ rad m}^{-2}$, which is not the case in this region near the Galactic anti-center (see below).

In Fig. 6.2 we show the polarization vectors around a few of the deepest ‘canals’, superimposed on grey scale plots of $P$, in two frequency bands. The area shown is indicated in Fig. 6.1. The polarization vectors on either side of the ‘canals’ are quite close to perpendicular, demonstrating that the ‘canals’ are produced by beam depolarization. This perpendicularity applies to all ‘canals’, irrespective of frequency band and is very convincing, especially because everywhere else the polarization vectors vary quite smoothly (if significantly!).

Beam depolarization creates ‘canals’ that are one beam wide, which is exactly what we observe. This implies that the $90^\circ$ ‘jump’ must occur on angular scales smaller than the beam width. At $\sim 2^\prime$ resolution (about twice that in Fig. 6.2), the ‘canals’ indeed seem unresolved, but the decrease in S/N-ratio precludes conclusions on even smaller scales (the original data have $0.8^\prime$ resolution).

Additional evidence that the ‘canals’ are due to beam depolarization is statistical.
Figure 6.2: Polarized intensity $P$ at 349 MHz (left) and 360 MHz (right) of the area inside the box in Fig. 6.1. Polarization angles and intensities are indicated by the vectors, which are sampled at locations $4'$ apart (independent beams). Note $\Delta \phi_{\text{pol}} \approx 90^\circ$ across low-$P$ 'canals'.

We defined 'canal-like' points from the observed values of $P$, as follows. For each point in the mosaic we compared the observed value of $P$ with the $P$-values in pairs of two diametrically opposed neighboring (adjacent) points. If the value of $P$ in the central point was less than a certain small fraction of the values in both comparison points, the point was defined 'canal-like'. This definition mimics the visual detection 'algorithm'.

In the top panel of Fig. 6.3 we show the distribution of the difference between the $\phi_{\text{pol}}$'s in the two adjacent points that define the 'canal-like' points, for a $P$-threshold of 30%. The $\Delta \phi_{\text{pol}}$-distribution peaks at $90^\circ$, fully consistent with the beam depolarization hypothesis. This conclusion is reinforced by a comparison with the distribution of $\Delta \phi_{\text{pol}}$ (again for diametrically opposed adjacent neighbors) of all points for which $P$ is between 1.0 and 2.0 times larger than both $P$-values in the two diametrically opposed neighboring points, which is shown in the bottom panel of the same figure.

Similar 'canals' were noted by Uyaniker et al. (1999) and Duncan et al. (1998), who also invoked beam depolarization. Yet, Fig. 6.3 is the first quantitative proof for this explanation.

6.4 The cause of the 'jumps' in polarization angle

Two processes can cause jumps in polarization angle $\phi_{\text{pol}}$ across the 'canals': a sudden change in $RM$ across the 'canals', and a jump in intrinsic $\phi_{\text{pol}}$ of the emission incident on the Faraday screen. A large change in intrinsic $\phi_{\text{pol}}$ implies a change in magnetic field direction and is therefore quite difficult to understand in view of the absence of structure in total intensity $I$ at the more than 2% level (see Sect. 6.2). On the other hand, variations in the $RM$ of the Faraday screen would seem to be quite natural, if
Figure 6.3: Top panel: polarization angle difference $\Delta \phi_{\text{pol}}$ between two points on opposite sides of a ‘canal-like’ point (see text for definition). A clear preference for $\Delta \phi_{\text{pol}} \approx 90^\circ$ across ‘canals’ is visible. Bottom panel: $\Delta \phi_{\text{pol}}$ between two points on opposite sides of non-canal-like points. For the non-canal-like points, $\langle \Delta \phi_{\text{pol}} \rangle \approx 0^\circ$, rather than $90^\circ$.

not unavoidable.

Discontinuities in $RM$ must play an important role in producing the ‘canals’, because the ‘canals’, although similar in adjacent frequency bands, generally do not occur in all bands, and certainly are not identical in the different bands (see Fig. 6.2). This indicates that the jumps in $\phi_{\text{pol}}$ are mainly due to changes in $RM$. However, the question is if the jumps in $\phi_{\text{pol}}$ are indeed accompanied by jumps in $RM$ of the right magnitude so that $\Delta \phi_{\text{pol}} = 90^\circ$ is produced at the frequency where the ‘canal’ is best visible.

In principle, the determination of $RM$ only involves a simple linear fit of the polarization angles in the five frequency bands (at 341, 349, 355, 360 and 375 MHz) vs. $\lambda^2$, but in practice several complications may arise. First, the observed values of $\phi_{\text{pol}}$ may be biased due to imaging effects (like offsets) in the Stokes $Q$- and $U$-maps from which $\phi_{\text{pol}}$ is derived (cf. Wieringa et al. 1993). Our data indicate that, in the maps of this region of sky, such offsets are quite small, so that the bias in the $\phi_{\text{pol}}$ values is small. Second, it is not obvious that the assumption of pure Faraday rotation ($\phi(\lambda) \propto \lambda^2$) is
supported by the data (see Chapter 4).

In Fig. 6.4 we show an array of plots of $\phi_{\text{pol}}(\lambda)$ vs. $\lambda^2$ for independent beams in the small region (indicated in Fig. 6.2) that contains two clear ‘canals’. As can be seen, a direct determination of $\Delta R M$ across the ‘canals’ is not at all trivial. Without knowing the position of the ‘canals’, one probably would have some trouble to find the ‘canals’ from discontinuities in $R M$ distribution alone, due to the uncertainties in the $R M$-estimates, which sometimes are considerable. On the other hand, if one knows where the canals are one can identify some related ‘jumps’ in $R M$.

From the present data, it seems quite likely that the ‘canals’ are primarily due to quite abrupt and relatively large changes of $R M$, with $\Delta R M/R M$ ranging from $\sim 0.3$ to more than 1 (at least in this region of sky). Note that in this region the $R M$s are in the range from $\sim 10$ to $+10$ rad m$^{-2}$ (also confirmed by several polarized extragalactic radio sources in these same observations). However, a more robust conclusion about the relation between $\Delta \phi_{\text{pol}}$ and $\Delta R M$ requires a detailed analysis of more, and more sensitive data, and a careful error analysis.
6.5 Implications for the structure of the local ISM

Because we have not yet reached a quantitative conclusion about the suspected correlation between $\Delta \phi_{\text{pol}}$ and $\Delta R.M.$, it is not possible to give a full discussion of the implications that these polarization data have for the small-scale structure of the warm ISM. However, the data discussed here show the great promise that high-resolution, multi-band polarization data hold for the study of the ISM, especially on small scales where pulsars and extragalactic radio sources cannot give much information.

Fortunately, more and more sensitive radio polarization data (in different regions of sky) are forthcoming. In addition, information must be obtained about the electron density in the warm ISM on the relevant scales (e.g. through H$\alpha$ measurements), as well as on the other components in the ISM (like e.g. the HI).

While we fully realize the preliminary nature of the conclusions presented, we feel justified to speculate somewhat on the possible implications of the ‘canals’. Structure in $R.M.$ reflects structure in $B_{||}$ and/or $n_e$ in the ISM. However, as the $R.M.$ is an integral over the entire line of sight, the large $\Delta R.M./R.M.$ values that are implied by our observations may give a very specific message. In particular, we consider it unlikely that the large $\Delta R.M./R.M.$ values are produced mainly by variations in electron density. Instead, they may be indicating a turbulent ISM with varying (reversing) magnetic field structures, as modeled in recent MHD simulations (see e.g. Mac Low & Ossenkopf 2000; Vázquez-Semadeni & Passot 1999).

Acknowledgements

The Westerbork Synthesis Radio Telescope is operated by the Netherlands Foundation for Research in Astronomy (NFRA) with financial support from the Netherlands Organization for scientific research (NWO). MH is supported by NWO grant 614-21-006.

References

de Bruyn A.G., Katgert P., 2000, in preparation
Polarimetric imaging at 325 MHz of a region in the WENSS survey at $137^\circ < l < 173^\circ$ and $-3^\circ < b < 30^\circ$

D. H. F. M. Schnitzeler, M. Haverkorn, P. Katgert, A. G. de Bruyn

Abstract

Polarization data from a part of the Westerbork Northern Sky Survey (WENSS) at 325 MHz are analyzed. Distinct structure is visible in polarization angle, mostly aligned with the Galactic plane. We estimate linear gradients of polarization angle in Galactic longitude and latitude. The gradients are computed in intervals, which are averaged to estimate the linear gradient across the field. We find that the polarization angle is approximately constant along Galactic longitude. However, the gradient of polarization angle in Galactic latitude is $\Delta \phi / \Delta b \approx -2.1$ radians per degree. If the gradient in rotation measure $\Delta M$ can be derived directly from the gradient in polarization angle, the gradient in $\Delta M$ is $\Delta M / \Delta b \approx -2.5$ m$^{-2}$ per degree, which is higher than previous estimates in this region. However, previous studies were done at lower resolution and higher frequency. Therefore, depolarization across the beam width is expected to be lower in our data, and depolarization along the line of sight higher, which may alter the $\Delta \phi$ and $\Delta M$ values. A high random component of the Galactic magnetic field may increase the depolarization along the line of sight significantly, which decreases the path length over which polarized radiation is visible. Therefore, a decrease in the relative strength of the random component of the Galactic magnetic field with increasing Galactic latitude would indicate an increase in path length. This could possibly explain the observed increase in $|RM|$ with Galactic latitude.

7.1 Introduction

Rengelink et al. (1997) carried out the Westerbork Northern Sky Survey (WENSS), which is a radio survey at 325 MHz, done with the Westerbork Synthesis Radio Telescope (WSRT) of the whole sky north of $\delta = 30^\circ$. The WENSS survey yielded all
Stokes parameters $I, Q, U$ and $V$, and although Rengelink et al. concentrated on extragalactic sources (and only on Stokes $I$), WENSS also contains a wealth of diffuse polarization data over a large region of sky.

The diffuse polarization exhibits abundant structure on many scales, both in polarization angle $\phi$ and in polarized intensity $P$, which is mostly unaccompanied by corresponding structure in total intensity. This indicates that most of the structure in polarization must be due to Faraday rotation and depolarization effects in the ISM. Therefore, diffuse polarization provides a unique tool in the study of properties of the ISM, in particular the large-scale (regular) and small-scale (random) components of the Galactic magnetic field. The WENSS survey is a single-frequency survey, so rotation measure $R\!M$ cannot be determined. However, differential $R\!M$ can be obtained through the variation in polarization angle $\phi$, as $\Delta R\!M = \Delta \phi / \lambda^2$, assuming a constant intrinsic polarization angle. On this assumption, single-frequency observations of the polarization angle can yield information on the structure in $R\!M$.

Study of the small-scale structure in polarization was initiated by Wieringa et al. (1993), who demonstrated the plausibility of Faraday rotation as a mechanism that creates the small-scale structure in polarization angle. Many studies of the small-scale structure in the diffuse polarized background have been made since then, all of them (except the studies described in this thesis) at frequencies of 1.4 GHz or higher. As such high frequencies probe high $R\!Ms$, most studies focused on the Galactic plane. Distinct objects such as supernova remnants or HII regions complicate the evaluation of the large-scale magnetic field in the plane, and information at intermediate and high Galactic latitudes is lacking. This is why the polarization data from the WENSS survey are so valuable: the WENSS polarization data extend over $137^\circ < l < 173^\circ$ and $-3^\circ < b < 30^\circ$, which gives an opportunity to study the structure of the ISM at high resolution over large ranges of Galactic longitude and latitude. In this paper, we study the change in polarization angle over this region, to be able to derive information on the large-scale characteristics of the Galactic magnetic field in the warm ISM.

Section 7.2 describes the observations and the correction for ionospheric Faraday rotation, and contains the observational results. In Section 7.3, we describe a procedure used to quantify the large-scale structure in polarization angle. Section 7.4 presents the large-scale gradients in polarization angle that we derived with this procedure, and in Section 7.5 we discuss a possible interpretation of the polarization angle gradients in terms of rotation measure. Finally, Section 7.6 presents our conclusions.

7.2 The observations

7.2.1 The WENSS survey

The Westerbork Northern Sky Survey (WENSS) is a low-frequency radio survey that covers the whole sky north of $\delta = 30^\circ$ at 325 MHz with a frequency bandwidth of 5 MHz, to a limiting flux density of approximately 18 mJy (5$\sigma$), and with a resolution of $54^\prime \times 54^\prime$ cosec $\delta$.\footnote{WENSS also includes a survey at 609 MHz over part of the WENSS region, which we will not discuss here.} To obtain a large field of view in an acceptable amount of time, the mosaicking technique was used, i.e. the telescope cycles through approximately...
80 pointings during a 12hr observation, integrating each frame for 20 seconds, with 10 seconds time to move the telescope to the next pointing position. This yields a homogeneous coverage of the \((u,v)\)-plane, and allows a large field of view to be observed in one 12hr observing session. Six 12hr syntheses were performed per mosaic with different telescope configurations, which effectively results in a smallest baseline of 36m, and a baseline increment of 12m. With six 12hr syntheses, the distance of the first grating ring from the pointing center is 4.4\(^\circ\) at 325 MHz.

The total area observed in 6 \(\times\) 12hr observations is over 100 square degrees, and its shape depends on declination. The WENSS consists of 68 of such mosaics in bands at declinations above 30\(^\circ\).

The data reduction and calibration of the WENSS survey was done with the NEWSTAR data reduction package, see Chapter 2. For calibration of the gains and phases of the linear dipoles, we used observations of the unpolarized calibrator sources 3C48 and 3C147. The absolute flux scale at 325 MHz is based on a value of 26.93 Jy for 3C286 (Baars et al. 1977).

### 7.2.2 The polarization mosaics

Measurements of diffuse polarization are vulnerable to polarized solar radiation, received through distant sidelobes. Therefore, only observations done (almost) completely at night can be used for polarimetry. We selected five adjacent mosaics that were observed largely or completely at night, the positions and outlines of which are shown in Fig. 7.1. Information about the five mosaics is presented in Table 7.1. The central declination of the two lower mosaics is 50\(^\circ\), and that of the three upper mosaics 66\(^\circ\). A Gaussian taper was applied to the data to enhance the sensitivity to extended emission. The Gaussian taper function has a value of 0.25 at a baseline of 250m, which yields a resolution of 6.7\(^\prime\) \(\times\) 6.7\(^\prime\) cosec \(\delta\).

Polarization calibration requires extra steps beyond the total intensity calibration.
<table>
<thead>
<tr>
<th>mosaic</th>
<th>WN50-J74</th>
<th>WN50-J90</th>
<th>WN66-J64</th>
<th>WN66-J83</th>
<th>WN66-J02</th>
</tr>
</thead>
<tbody>
<tr>
<td>(\alpha, \delta)</td>
<td>((74^\circ, 50^\prime))</td>
<td>((90^\circ, 50^\prime))</td>
<td>((64^\circ, 66^\prime))</td>
<td>((83^\circ, 66^\prime))</td>
<td>((102^\circ, 66^\prime))</td>
</tr>
<tr>
<td>(\nu) (MHz)</td>
<td>325</td>
<td>325</td>
<td>325</td>
<td>325</td>
<td>325</td>
</tr>
<tr>
<td>resolution</td>
<td>(6.7 \times 8.7^\prime)</td>
<td>(6.7 \times 8.7^\prime)</td>
<td>(6.7 \times 7.3^\prime)</td>
<td>(6.7 \times 7.3^\prime)</td>
<td>(6.7 \times 7.3^\prime)</td>
</tr>
<tr>
<td>date</td>
<td>36m</td>
<td>92/01/07</td>
<td>92/01/13</td>
<td>92/01/12</td>
<td>92/01/09</td>
</tr>
<tr>
<td></td>
<td>48m</td>
<td>91/12/31</td>
<td>92/01/06</td>
<td>92/01/05</td>
<td>92/01/02</td>
</tr>
<tr>
<td></td>
<td>60m</td>
<td>91/12/24</td>
<td>91/12/30</td>
<td>93/01/10</td>
<td>91/12/26</td>
</tr>
<tr>
<td></td>
<td>72m</td>
<td>91/12/06</td>
<td>91/12/02</td>
<td>91/12/01</td>
<td>91/12/05</td>
</tr>
<tr>
<td></td>
<td>84m</td>
<td>91/12/10</td>
<td>91/12/16</td>
<td>91/12/15</td>
<td>91/12/12</td>
</tr>
<tr>
<td></td>
<td>96m</td>
<td>91/12/17</td>
<td>91/12/23</td>
<td>91/12/22</td>
<td>91/12/19</td>
</tr>
<tr>
<td>start time</td>
<td>36m</td>
<td>15:24</td>
<td>16:06</td>
<td>14:26</td>
<td>17:36</td>
</tr>
<tr>
<td>(UT)</td>
<td>48m</td>
<td>15:52</td>
<td>16:33</td>
<td>14:53</td>
<td>16:21</td>
</tr>
<tr>
<td></td>
<td>60m</td>
<td>16:20</td>
<td>17:01</td>
<td>14:31</td>
<td>16:49</td>
</tr>
<tr>
<td></td>
<td>72m</td>
<td>17:41</td>
<td>18:51</td>
<td>17:11</td>
<td>18:12</td>
</tr>
<tr>
<td></td>
<td>84m</td>
<td>17:48</td>
<td>17:56</td>
<td>16:16</td>
<td>17:44</td>
</tr>
<tr>
<td></td>
<td>96m</td>
<td>16:47</td>
<td>17:28</td>
<td>15:48</td>
<td>17:16</td>
</tr>
</tbody>
</table>

Table 7.1: Observational details of the five WENSS mosaics discussed in this paper. We list the central right ascension and declination, the observing frequency \(\nu\), the resolution (for the center of the mosaic), the observing dates and start times of the observation in UT. The latter two are given for the individual 12hr syntheses, with minimum baseline increments (spacings) from 36m to 96m. The 36m-spacing is observed twice for WN66-J83, and parts of both observations are used in the construction of the mosaic.

First of all, unpolarized calibrators were used to determine dipole and ellipticity errors in the two orthogonal dipole feeds (see Chapter 2). A minor amount of signal leakage from Stokes \(U\) to Stokes \(V\) was visible. This is due to a phase-zero difference between the signals in the \(XY\) and \(YX\) visibilities (see Chapter 2). It was corrected by manually adding a small phase difference of 9° between the \(XY\) and \(YX\) visibilities.

At these low frequencies, the Faraday rotation due to the ionosphere becomes important. Because we can only determine differential \(R.M\), we only need to correct for changes in the ionospheric Faraday rotation during the six observing sessions for one mosaic. The differential correction between different 12hr syntheses was derived from the extragalactic point sources in the mosaics. If the sources are bright enough, their polarization angle can be determined in each of the six 12hr observing sessions separately. Polarization angle differences between the observing sessions are solely caused by differences in ionospheric Faraday rotation. The polarization angle data in each 12hr observation was corrected for this angle difference before combining the six 12hr periods to obtain the final mosaic. The ionospheric Faraday rotation corrections vary from 0.7° to 17°, and have an average of about 5°. This low value indicates that the ionospheric Faraday rotation was very stable during the two-month period in which the observations were taken.

Changes in ionospheric Faraday rotation during one 12hr synthesis cannot be accounted for. However, a large change in Faraday rotation during one 12hr observation
would give a variable polarization angle during the night, and the polarized signal integrated over the 12 hours would be largely depolarized. The fact that polarized extragalactic point sources are still visible, even if only one 12hr observing session is considered, is an indication that differential Faraday rotation during one 12hr period is not very large.

Off-axis instrumental polarization is largely eliminated in the mosaicking technique, resulting in an instrumental polarization of less than 1% in the field. Only at the edges of the field, off-axis instrumental polarization can reach very high values, therefore all data closer than \( \sim 0.8^\circ \) from the outer edges are not included in the analysis.

### 7.2.3 Combination of the mosaics into one “supermosaic”

To obtain a large field of view with a size of about \( 30^\circ \times 35^\circ \), the five mosaics must be combined into one “supermosaic”. But after the correction for ionospheric Faraday rotation discussed above, each mosaic still contains an arbitrary polarization angle zero-point. To make the normalization in polarization angle consistent in all mosaics, we compared polarization angles in overlap regions between two mosaics. These overlap regions are a few degrees-wide, see Fig. 7.1.

The difference in polarization angle for the same position in the two mosaics \( i \) and \( j \) is \( \Delta \phi_{ij} = \phi_i(x) - \phi_j(x) \) for position \( x \). Only positions where the polarized intensity was above a threshold of 15 mJy/beam were used, to make certain that the polarization angle at that position is well-defined. \( \Delta \phi_{ij} \) was computed for the five largest overlapping regions, which yielded angle corrections between 2° and 8°. Note that the consistency between the mosaics is inherent to the observing strategy of observing all mosaics in a particular telescope configuration on successive nights. We checked that the values for \( \Delta \phi_{ij} \) were consistent by calculating “closure errors”, which were found to be \( < 1^\circ \). Taking the mosaic WN50,000 as a reference, we applied the values of \( \Delta \phi_{ij} \) to the other four mosaics, to correct the polarization angle.

The supermosaic was constructed using the reduction package NEWSTAR and was regridded from equatorial coordinates into Galactic coordinates before the analysis. This yields polarization data in an irregularly shaped region within \( 137^\circ < l < 173^\circ \) and \( -3^\circ < b < 30^\circ \).

### 7.2.4 The observational results

From the Stokes \( Q \) and \( U \) maps, polarized intensity \( P \) and polarization angle \( \phi \) are derived, which are shown in Figs. 7.2 and 7.3.

The map of polarized intensity shown in Fig. 7.2 is saturated at an intensity of 150 mJy/beam, which is equivalent to a polarized brightness temperature of 10 - 15 K, depending on declination. (The conversion factor between intensity and brightness temperature is 100 mJy/beam = 10.3 \( \sin \delta \) K at 325 MHz for a beam of \( 6.7^\prime \times 6.7^\prime \) csc \( \delta \).) Higher polarized intensities are mostly due to high instrumental polarization at the edges of the field. The filamentary structure of high polarized intensity around \( (l, b) = (161^\circ, 15^\circ) \) was reobserved at multiple frequencies, the results of which are described in detail in Chapter 4.
Figure 7.2: Polarized intensity $P$ in the supermosaic, made out of 5 mosaics of the WENSS survey at 325 MHz. The data is tapered with a Gaussian taper with a value of 0.25 at a baseline of 250m, which yields a resolution of $\approx 6.7\text{arcmin}$, and is regridded to Galactic coordinates. The maximum $P$ shown in white is 150 mJy/beam, which is equivalent to a polarized brightness temperature of 10 - 15 K, depending on declination.

The map of polarization angle is presented in Fig. 7.3, with coding from $-90^\circ$ (black) to $90^\circ$ (white), so that white and black denote the same angle. The polarization angles appear to be coherent over large scales, in the direction of Galactic longitude as well as latitude. Furthermore, at all positions where $P < 9.1$ mJy/beam ($1\sigma$), the signal-to-noise ratio was so low that the error in the polarization angle is too high to allow reliable determination of the angle. Therefore, these data were discarded as well.

In both the polarized intensity and the polarization angle, there is a change in the scale of the structure at $b \approx 20^\circ$. This could be related to a region of high and regular polarization (the “fan region”), that extends over $160^\circ \leq l \leq 120^\circ$ and $0^\circ \leq b \leq 20^\circ$ (Westerhout et al. 1962, Bronw and Spodstra 1976). However, it cannot be concluded that high polarized intensity is not present at high Galactic latitudes, because previous
Figure 7.3: Polarization angle $\phi$ in the supermosaic from the WENSS survey at 325 MHz in Galactic coordinates, with the same taper as in Fig. 7.2. The polarization angle range is $[-90^\circ, 90^\circ]$, so that white denotes the same angle as black.

Low-frequency polarimetric observations at high Galactic latitudes with the WSRT also show abundant structure in polarized intensity and polarization angle (Kagt and de Bruyn 1999). More diffuse polarization observations at high Galactic latitudes and at low frequencies are needed to investigate this apparent change in structure.

### 7.3 Determination of the gradients in polarization angle

The most remarkable feature in Fig. 7.3 is the coherence of polarization angle over large parts of the WENSS polarization field. On the assumption that polarization angle has a one-to-one relation with rotation measure $RM$, this allows us to study the large-scale characteristics of $RM$ by quantifying the large-scale gradient of polarization angle, e.g. along Galactic longitude and Galactic latitude.

However, much small-scale structure is present as well in the polarization angle. In
principle, the structure in polarization angle could be described fully by power spectra or structure functions in Galactic longitude and latitude (as was done in Chapter 8). However, on large angular scales, the n180° ambiguity of the polarization angle is a problem that cannot be easily remedied.

Therefore, we constructed a simple scheme to attempt to separate small-scale and large-scale structure in polarization angle, which is illustrated in Fig. 7.4. We sampled the data in bins of size $\theta_l \times \theta_b$ in the directions of Galactic longitude and latitude, respectively. The polarization angles within a bin are averaged to one value $\phi_{\text{bin}}$, where the n180° ambiguity of the polarization angle was taken into account. The error in the average polarization angle is the error in the mean $\sigma_{\phi_{\text{bin}}} = \sigma / \sqrt{N}$, where $\sigma$ is the standard deviation of the $\phi$ distribution in a bin, and $N$ is the number of independent beams in a bin. Only values of $\phi_{\text{bin}}$ for which the following conditions hold are included in the analysis: (1) the standard deviation in $\phi_{\text{bin}}$ must be smaller than 20°; (2) more than 50% of the beams in a bin has to be included, according to the constraints mentioned in Section 7.2; and (3) the number of usable beams in a bin must be $\geq 1$. The latter restriction is only important for very small bin sizes.

Our first goal is to determine the large-scale gradient in $\phi$, so we make a linear
fit to the polarization angle as a function of position. However, determination of the gradient in the angle is hampered by the ambiguity in the angle. This is illustrated in Fig. 7.5, which shows $\phi_{bin}$ as a function of Galactic latitude for three adjacent strips at $l \approx 147.6^\circ$. The ambiguity in $\phi_{bin}$ was resolved (for each strip independently) by minimizing $\phi_{bin,i} - \phi_{bin,j}$, where $i$ and $j$ denote two adjacent usable bins. The polarization angle is similar for the three strips, but shows differences in angle of $\pm 180^\circ$ between strips. By adding $\pm 180^\circ$ to strips, the angle determinations could be made almost the same for the three strips. However, it is impossible to decide which angle gradient is the correct one without external constraints.

We tried to determine the best polarization angle gradient by assuming a range of linear gradients $\phi = ax + \phi_0$ and determine the best fit to the data. To estimate the best-fit linear gradient, we rotated the polarization angle values by n $180^\circ$ where $n$ was such that the corrected values were closest to the imposed gradient value $\phi = ax + \phi_0$. We then computed a reduced $\chi^2$ of the data set as

$$\chi^2_{red} = \frac{1}{N-2} \sum_{i=1}^{N} \frac{(\phi_i - (ax + \phi_0))_i^2}{\sigma_{bin}^2}$$

For a range of gradients $a$ and polarization angle zero-points $\phi_0$, the $\chi^2_{red}$ values were computed, and the gradient with the lowest $\chi^2_{red}$ value was adopted as most likely. The range of gradients was $-25 < a < 25^\circ$/beam, with an increment of $0.5^\circ$/beam. The values of polarization angle zero-points spanned the entire range of $-90^\circ < \phi_0 < 90^\circ$ with an increment of $0.1^\circ$. In this way, we derived the best linear fit to the polarization angle as a function of position, despite the $n 180^\circ$ ambiguity in the polarization angle. Only linear fits in which the reduced $\chi^2$ was smaller than a maximum value $\chi^2_{max}$ were included in the analysis.

When the linear gradient in $\phi_{bin}$ was computed over the entire length of the field, we did not find slopes $a$ and zero-points $\phi_0$, which clearly gave a better linear fit than other values. This indicates that the data show too much small-scale structure to allow the determination of a large-scale gradient across the entire length of the field.
Free parameters in large-scale gradient procedure

<table>
<thead>
<tr>
<th>Parameter</th>
<th>Description</th>
</tr>
</thead>
<tbody>
<tr>
<td>$\theta_t, \theta_b$</td>
<td>bin size</td>
</tr>
<tr>
<td>$n_{int}$</td>
<td>interval length in bins</td>
</tr>
<tr>
<td>$N_{b_{\text{min}}}$</td>
<td>minimum number of data points in the set of $\phi_{\text{bin}}$ in one interval</td>
</tr>
<tr>
<td>$\chi^2_{\text{max}}$</td>
<td>maximum reduced $\chi^2$ of linear fit of $\phi_{\text{bin}}$ as a function of position</td>
</tr>
<tr>
<td>$N_{l_{\text{min}}}$</td>
<td>minimum number of data points in the set of slope determinations $a_j$</td>
</tr>
</tbody>
</table>

Table 7.2: Free parameters in the procedure to determine the large-scale gradient in polarization angle along Galactic longitude and latitude.

Therefore, we divided each strip into intervals, which overlapped to yield Nyquist sampling. In each interval separately, a linear fit of $\phi_{\text{bin}}$ as a function of position was constructed in the way described above. This yields a number of evaluated slopes equal to the number of intervals within a strip. The distribution of these slopes yields in most cases a distribution with a clear maximum, and we adopt the average slope as the best estimate for the large scale gradient in $\phi$. If the number of reliably determined $\phi_{\text{bin}}$ values in an interval was smaller than a minimum $N_{b_{\text{min}}}$, the interval was ignored. Clearly, using fixed length intervals, we only probe a single angular scale. By varying interval lengths one obtains information on the spatial scales present in the polarization angle.

In some strips in Galactic latitude, the distribution of slopes per interval was clearly bimodal. In these cases, a bimodality filter was applied, in which the part of the distribution containing the least number of points was discarded (see Section 7.4.2).

This procedure results in a distribution of polarization angle gradients $a_j$ for intervals $j$ along one strip. If the number of data points (allowed intervals) in this set of $a_j$ values is higher than a minimum value $N_{l_{\text{min}}}$, the average slope $a_s$ of the strip was determined, with $\sigma_{a_s}$ the error in the mean of the distribution.

### 7.3.1 Optimization of parameters

The free parameters that can be optimized in the procedure are given in Table 7.2.

Optimization of the dimensions of the bins $\theta_t$ and $\theta_b$ was done within a range of 1 to 10 beams. The bin height $\theta_t$ cannot be too large because the gradient in the polarization angle in Galactic latitude is high (see Fig. 7.3). A large bin height $\theta_b$ yields a large range in polarization angle, so that a reliable average cannot be determined. On the other hand, the large-scale angle structure in longitude varies only very slowly, so that larger bins in the $l$-direction can be taken. We found that with $\theta_t \times \theta_b = 2 \times 2$ beams for strips in longitude and $\theta_t \times \theta_b = 4 \times 1$ beams for strips in latitude the small- and large-scale components could best be separated.

The interval length $n_{int} = 6$ bins in $l$, and $n_{int} = 12$ bins in $b$. Shorter intervals in many cases do not contain enough bins with good data to determine a reliable gradient in polarization angle for the interval. The minimum number of data points in one interval and in the set of slope determinations were $N_{l_{\text{min}}} = 6$ and $N_{b_{\text{min}}} = 6$, respectively, to allow reliable determination of slope $a_j$. The maximum allowed reduced $\chi^2$ is varied. Note that in this analysis, $\chi^2$ must be interpreted as a measure of substructure, rather than of the quality of the data.
7.4.1 Polarization angle gradient along Galactic longitude

The results are displayed as average slopes $\alpha$ in a strip as a function of the coordinates of the strip. Fig. 7.6 presents the average slope in polarization angle for strips in $\ell$, as a function of $b$. The plot is made using a linear fit of polarization angle against position in one interval, viz., $X^{0.2}_{\text{red}} = 10$ (solid) and $X^{0.2}_{\text{red}} = 20$ (dashed). The two lines denote two cut-off values of the reduced $X^2$ in the linear fit of $\alpha$ for $b > 17\deg$. The slope $\gamma$ in $\ell$ is generally consistent with zero, but could be slightly negative. The slope $\gamma$ in $b$ is generally consistent with zero, but could be slightly negative. The slope $\gamma$ in $b$ is generally consistent with zero, but could be slightly negative.

The total region used in the analysis is rectangular. Strips in longitude are taken as the number of intervals which are able to determine the average slope for the $X^2_{\text{red}} < 10$ data.

7.4.2 Polarization angle gradient along Galactic latitude

The plot of the average slope per strip in $b$ as a function of $b$ is given in the top panel of Fig. 7.7. In $b$ the slope is distinctly non-zero, as expected from Fig. 7.3.

Figure 7.6: Average slopes in polarization angle along strips in Galactic longitude, as a function of Galactic latitude. The results are for the values of $X^2_{\text{red}}$, viz., $X^2_{\text{red}} = 10$ (solid line) and $X^2_{\text{red}} = 20$ (dashed line). The number next to the data points denote the number of weighted mean of the slopes determined for $X^2_{\text{red}} = 10$. The total region used in the analysis is rectangular. Strips in longitude are taken as the number of intervals which are able to determine the average slope for the $X^2_{\text{red}} < 10$ data.
Figure 7.7: Average slopes in polarization angle along strips in Galactic latitude, as a function of Galactic longitude. Notation as in Fig. 7.6. Top: average slope per interval determined from the complete distribution \( \alpha \). Bottom: average slope per interval determined after discarding the least occupied mode of bimodal distributions.

However, the errors in the average slopes are very large, indicative of a broad distribution of slopes within a strip. In fact, most of the strips in Galactic latitude show a bimodal distribution of slopes. This is shown in Fig. 7.8. The left-hand panels give the distribution of slopes, the right-hand panels show the intervals with negative and positive average slopes in \((l, b)\). The field is divided in strips in Galactic latitude, separated by the thick black lines. Each strip is divided in intervals (“bricks”), and two columns of intervals within one strip indicate the Nyquist sampling of the intervals. An interval is shown grey if the data was not usable, i.e. either there was no data in the interval, there were not enough bins with well-determined \( \phi_{\min} \) in the interval, or the reduced \( \chi^2 > \chi^2_{\text{max}} \). A black “brick” denotes an interval with a negative slope, and a white “brick” an interval with a positive slope. The top panels are for \( \chi^2_{\text{max}} = 5 \), the bottom panels for \( \chi^2_{\text{max}} = 60 \), which includes all linear fits.

The figure clearly shows a bimodal distribution of slopes. The intervals with positive slopes (about 1 in 5) are distributed quite regularly over the field. Exclusion of intervals
Figure 7.8: Left: histograms of the average slope in polarization angle per interval, for the complete field. Right: layout of intervals in strips along $b$. Strips are divided by thick black lines, each "brick" is an interval, and the Nyquist sampling is shown by means of two columns of intervals within each strip. Grey "bricks" indicate that no data is available or usable, black "bricks" denote negative slopes and white ones positive slopes. The top figure shows $\chi^2_{max} = 5$, the bottom figure $\chi^2_{max} = 60$.

with positive slopes in the estimate of the average slope results in the bottom panel of Fig. 7.7. The errors in the slope determination are much lower, and the average is consistent with a constant slope of $\sim -2.1$ radians per degree.

The intervals at the edges of the latitude range contain less usable intervals than the central range of $5^\circ < b < 23^\circ$. For $b > 23^\circ$, the character of structure in polarization angle changes, as is visible in Fig 7.3. Small-scale structure in polarization angle is very prominent. If the scale of the structure in polarization angle is comparable to or smaller than a bin size, the determination of the average $\phi_{bin}$ will have a large error, and the bin is considered not valid. Below $b = 5^\circ$, there is no data available in some positions, and too much small-scale structure in others.

Over a large range in $\chi^2_{max}$, the characteristics of the distribution of slopes hardly change. The ratio of positive and negative slopes is independent of $\chi^2_{max}$, and the shape of the histogram of slopes does not change with increasing $\chi^2_{max}$ either. This indicates that the average slope of $\sim -2.1$ radians per degree is characteristic for the region, although small-scale structure is non-negligible and is apparent from an increase in the
\( \chi^2 \) values of the linear fits.

At present, we probe only the spatial scale of the length of the interval, viz. \( \sim 1.4^\circ \). Fig. 7.8 indicates that the negative gradient in polarization angle is present on spatial scales larger than this, as large as \( \sim 5^\circ - 10^\circ \).

### 7.5 Large-scale differential rotation measure

As the WENSS polarization region contains only single-frequency data, rotation measure \( RM \) cannot be determined. However, the gradient in \( RM \) can be estimated from the gradient in polarization angle \( \Delta \phi / \Delta b \).

In the case of pure Faraday rotation, the polarization angle is \( \phi = \phi_0 + RM \lambda^2 \). If \( \phi_0 = \) constant, the gradient in \( RM \) is directly proportional to the gradient in polarization angle \( \Delta RM / \Delta b = \Delta \phi / \Delta b / \lambda^2 \).

The assumption that \( \phi_0 = \) constant is probably reasonably good. In these diffuse polarization observations, and many others (among those discussed in Chapters 4 and 5), there is abundant structure in polarization angle and polarized intensity, but no corresponding structure in total intensity \( I \). This indicates that the synchrotron emission must be uniform, while the structure in polarization is created by Faraday rotation and depolarization in the magneto-ionic medium. A uniform synchrotron emission is not consistent with structure in the intrinsic polarization angle \( \phi_0 \), so that probably \( \phi_0 = \) constant is a good approximation.

The assumption of the structure being due to pure Faraday rotation is most likely not valid (Chapter 3). Therefore, depolarization mechanisms would be important and can modify polarization angles so that \( \phi \neq \phi_0 + RM \lambda^2 \), and \( \Delta \phi / \Delta b / \lambda^2 \neq \Delta RM / \Delta b \).

Although random modification of polarization angle due to depolarization on scales of the bin will be averaged out in the averaging of polarization angles within a bin, depolarization which is uniform on larger scales is less easily distinguishable, and will complicate the derivation of slopes in intervals. However, the ordered appearance of polarization angle in Fig. 7.3 on large scales suggests that modification of polarization angles due to depolarization is not dominant. We will adopt the first order approximation that \( \Delta \phi / \Delta b \propto \Delta RM / \Delta b \) here. In this case, the gradient in \( RM \) in the direction of Galactic latitude is \( \Delta RM / \Delta b \approx -2.5 \) rad m\(^{-2}\) per degree.

The earliest \( RM \) determinations from diffuse polarization data in the direction of the WENSS polarization region were made by Bingham and Shakeshaft (1967). Using four radio surveys at frequencies from 408 MHz to 1407 MHz, they derived rotation measures among others in the area \( 170^\circ > l > 130^\circ \) and \( 0^\circ < b < 20^\circ \) with a resolution of \( \sim 2^\circ - 3^\circ \). They computed a \( RM \) map which shows a gradient approximately along Galactic latitude, where \( RM \)s range from \( RM \approx 7 \) rad m\(^{-2}\) at \( b = 3^\circ \) to \( RM \approx -8 \) rad m\(^{-2}\) at \( b = 20^\circ \). This indicates a gradient in \( RM \) of \( \sim -1 \) rad m\(^{-2}\) per degree. The change in sign of \( RM \) occurs approximately at Galactic latitude \( b = 10^\circ \).

From several radio surveys discussed in Brow and Spoedst (1976), Spoedst (1984) derived rotation measure values in a large part of the northern sky. The surveys with frequencies from 408 MHz to 1411 MHz have varying resolution of \( 2.3^\circ \) at 408 MHz to \( 0.6^\circ \) at 1411 MHz. Due to undersampling of the data, proper smoothing to the lowest resolution was not possible. Therefore, only structure in the \( RM \) determinations on
scales $\gtrsim 3^\circ$ should be trusted (Spoelstra 1984). The $R M$ observed by Spoelstra in the direction of the WENSS polarization region shows a gradient from $R M \approx 3$ rad m$^{-2}$ at $b = 0^\circ$ to negative $R M$s varying from $-6$ rad m$^{-2}$ to $-15$ rad m$^{-2}$ at $b = 20^\circ$. Above $b = 20^\circ$ $R M$ increases rapidly to $R M \approx 9$ rad m$^{-2}$ at $b = 27^\circ$. Therefore, the largest gradient obtained by Spoelstra is also $\sim -1$ rad m$^{-2}$ per degree in the region of $b \leq 20^\circ$, with a sign change around $b = 10^\circ$. This result is thus consistent with that of Bingham and Sparks. However, both $R M$ gradients are more than twice as low as $\Delta R M/\Delta b$ estimated from the present data.

$R M$ determinations from nearby pulsars show predominantly negative $R M$s in the range $R M \approx -20$ to $-40$ rad m$^{-2}$ in the Galactic plane at $l = 120^\circ - 160^\circ$ (Rand and Kulkarni 1988, Rand and Lyne 1994). A few pulsars with positive $R M \approx 10$ rad m$^{-2}$ at $b \approx 7^\circ$ are located at a distance $< 3$ kpc. Polarized extragalactic point sources also show mainly negative $R M$s in the Galactic plane, with $R M \approx -20$ rad m$^{-2}$ at $l = 170^\circ$ to $R M \approx -150$ rad m$^{-2}$ at $l = 130^\circ$ (Simard-Normandin and Kronberg 1980). Above the plane, both negative and positive $R M$s are observed of order 10 - 30 rad m$^{-2}$.

Extragalactic point sources probe the complete line of sight through the Galaxy, whereas diffuse radiation which originates at large distances is probably mostly depolarized in passage through the magneto-ionic medium in the Galactic disk. The $R M$s from the diffuse radio polarimetric observations probably only probe the nearest few hundred parsecs at these low frequencies. Furthermore, extragalactic sources could have a $R M$ component intrinsic to the source or a halo around the source as well. Therefore, the $R M$ determinations from diffuse radiation can differ from $R M$s derived from extragalactic sources or distant pulsars.

The $R M$ determinations through diffuse polarization show a change in sign of $R M$ at approximately constant Galactic latitude around $b = 10^\circ$, which indicates a reversal of the Galactic magnetic field. As the diffuse polarization observations only probe the nearest part of the ISM, we cannot deduce whether this is a local reversal or not. However, the $R M$ data from extragalactic polarized sources and pulsars weakly suggest that the reversal persists over the complete line of sight through the Galaxy. On the plane of the sky, it seems to be present on scales of several tens of degrees.

The large gradient in $R M$ that we observe is caused by a combination of gradients, viz. that in the thermal electron density, in the length of the visible part of the line of sight, and in the parallel component of the magnetic field. The change of sign of $R M$ indicates that the parallel magnetic field component must change direction.

The path length through the medium is largely determined by the amount of depolarization along the line of sight in the medium, which yields a “depolarization horizon”. This depolarization horizon is frequency dependent, and is located at a shorter distance if there is more depolarization. In the Galactic plane, the ratio of random and regular magnetic field components $B_{r m}/B_{r g}$ is likely to be higher than out of the plane (Han and Qiao 1994, Indrani and Deshpande 1998). If so, the depolarization horizon would be nearer in the plane, so that the path length in the plane is smaller than at higher latitudes. This can cause an additional gradient in $R M$.

In Chapter 3, we constructed a depolarization model which was compared to two multi-frequency radio polarimetric data sets, discussed in Chapters 4 and 5. This model yielded an estimate of a polarization horizon which is shown in Fig. 7.9. This figure shows the observed fraction of polarization originating at a certain distance, as
Figure 7.9: Observed fraction of polarization originating at a certain distance as a function of that distance, for model parameters that best reproduce the observables from the two multi-frequency data sets discussed in Chapters 4 and 5 (solid lines). The dashed and dotted lines give the situation in which the random magnetic field is 10% and 50% higher than in the situation for the solid lines.

a function of that distance. The solid lines denote the model with parameters which approximately reproduce the results from the two multi-frequency observations. If the random magnetic field is enhanced by 10%, the observed fraction of polarization is given by the dashed lines, and an increase of the random magnetic field of 50% is denoted by the dotted lines. From the figure, we conclude that the polarization horizon is greatly influenced by \( B_{\text{ran}} / B_{\text{reg}} \).

If \( B_{\text{ran}} / B_{\text{reg}} \) is 10% higher in the plane than at higher latitudes of \( b \approx 20^\circ \), the distance of the polarization horizon can decrease by more than a factor 2 according to the model. For a modest change in Galactic magnetic field of \( B_\parallel = 0.5 \mu \text{G} \) at \( b = 0^\circ \) to \( B_\parallel = -0.5 \mu \text{G} \) at \( b = 20^\circ \), and assuming a constant \( n_e = 0.08 \text{ cm}^{-3} \), a change in path length from 600 pc at \( b = 20^\circ \) to 300 pc at \( b = 0^\circ \) can cause a change in \( R M \) from \( R M = 10 \text{ rad m}^{-2} \) at \( b = 0^\circ \) to \( R M = -20 \text{ rad m}^{-2} \) at \( b = 20^\circ \). We emphasize that this estimate uses many assumptions and is highly uncertain. However, it does show that a large gradient in \( R M \) can be obtained from a moderate decrease in the random component of the magnetic field towards higher Galactic latitudes.

In the \( 9^\circ \times 11^\circ \) region centered at \( (l, b) = (161^\circ, 16^\circ) \), discussed in Chapter 4, a gradient in \( R M \) was observed in the direction of Galactic longitude, i.e. opposite to the gradient in polarization angle found here. However, this gradient is only \( \sim 1 \text{ rad m}^{-2} \) per degree, and, as is already apparent from Fig. 7.3, this region is small enough with respect to the WENSS polarization region that the effect of the deviating gradient is averaged out.
7.6 Conclusions

Low-frequency polarimetric observations of large fields at non-zero latitudes represent a powerful tool to deduce information on the large-scale structure of the nearby warm ISM.

The observed polarization angle of the diffuse polarized radio background clearly shows spatial structure on scales of tens of degrees. The polarization angle is approximately constant in Galactic longitude and shows a large gradient in Galactic latitude. Small-scale structure is superimposed on this large-scale gradient.

The large-scale gradient in polarization angle in the directions of Galactic longitude or Galactic latitude can be estimated by evaluation of an average angle gradient in narrow strips along longitude or latitude. Due to the $\pm 180^\circ$ ambiguity of the polarization angle, the gradient was not determined for a complete strip. Therefore, the strips were divided in intervals, and the gradient of polarization angle in a strip was taken to be the average of the gradients in all intervals in the strip.

There is no significant polarization angle gradient in Galactic longitude. In Galactic latitude, the polarization angle gradient is distinctly non-zero, viz. $\Delta \phi / \Delta b = -2.1$ radians per degree.

Assuming that the intrinsic polarization angle is constant and that depolarization effects do not modify the polarization angle significantly, the gradient in polarization angle can be interpreted as a gradient in rotation measure $RM$. This gradient is observed in other observations of the diffuse radio polarization as well, although its magnitude is approximately twice as low in those studies (Bingham and Shakeshaft 1967, Spoelstra 1984). Because previous measurements were done at much lower resolution, beam depolarization is likely to be higher than in our data. Furthermore, the data were taken at higher frequencies than 325 MHz, so that in our data depolarization along the line of sight is likely to be more dominant.

An increase in the distance of the polarization horizon due to a decrease in the random magnetic field component with increasing Galactic latitude can yield large $RM$ gradients, which could explain the observed gradient in polarization angle.

Acknowledgements

The Westerbork Synthesis Radio Telescope is operated by the Netherlands Foundation for Research in Astronomy (ASTRON) with financial support from the Netherlands Organization for Scientific Research (NWO). The Wisconsin H-Alpha Mapper is funded by the National Science Foundation. MH is supported by NWO grant 614-21-006.

References

Indrani C., & Deshpande A. A., 1998, NewA 4, 331
Westerhout G., Soeger Ch. L., Brouw W. N., & Timmer J., 1962, BAN 16, 187
8

Characteristics of the structure in the Galactic polarized radio background at 350 MHz

M. Haverkorn, P. Katgert and A. G. de Bruyn, submitted to A&A

Abstract

Angular power spectra and structure functions of the Stokes parameters $Q$ and $U$, and polarized intensity $P$ are derived from three sets of radio polarimetric observations. Two of the observed fields have been studied at multiple frequencies, allowing determination of power spectra and structure functions of rotation measure $RM$ as well. The third field extends over a large part of the northern sky, so that the variation of the power spectra over Galactic latitude and longitude can be studied. The power spectra of $Q$ and $U$ are steeper than those of $P$, probably because a foreground Faraday screen creates extra structure in $Q$ and $U$, but not in $P$. The extra structure in $Q$ and $U$ occurs on large scales, and therefore causes a steeper spectrum. The derived slope of the power spectrum of $P$ is the multipole spectral index $\alpha_P$, and is consistent with earlier estimates. The multipole spectral index $\alpha_P$ decreases with Galactic latitude (i.e. the spectrum becomes flatter), but is consistent with a constant value over Galactic longitude. Power spectra of the rotation measure $RM$ show a spectral index $\alpha_{RM} \approx 1$, while the structure function of $RM$ is approximately flat. The structure function is flatter than earlier estimates from polarized extragalactic sources, which could be due to the fact that extragalactic source $RM$ probe the complete line of sight through the Galaxy, whereas as a result of depolarization, diffuse radio polarization only probes the nearby ISM.

8.1 Introduction

The warm ionized gaseous component of the Galactic interstellar medium (ISM) shows structure in density and velocity on scales from AU to several kiloparsecs. The Galactic magnetic field is coupled to the motions of the gas, and has an energy density comparable to that of the warm gas, so that gas and magnetic field are in complex interaction. Detailed knowledge of the turbulent nature of the warm ISM and the structure in
the Galactic magnetic field is essential for several fundamental studies of the Galaxy, including modeling of molecular clouds (e.g. Vázquez-Semadeni and Passot 1999, Ostriker et al. 2001), heating of the ISM (Minter and Balser 1997), star formation (e.g. Fernière 2001), and cosmic ray propagation (Chevalier and Fransson 1984).

Small-scale structure in the warm ISM and magnetic field can be well studied using polarimetric observations, of the radio synchrotron background in the Milky Way (e.g. Brouw and Spoelstra 1976, Wieringa et al. 1993, Duncan et al. 1997 and 1999, Uyaniker et al. 1999, Landecker et al. 2001, Gaensler et al. 2001), of pulsars (e.g. Rand and Kulkarni 1989, Ohno and Shibata 1993, Rand and Lyne 1994, Han et al. 1999) or of polarized extragalactic point sources (e.g. Simard-Normandin and Kronberg 1980, Clegg et al. 1992). At short wavelengths ($\lambda \lesssim 6$ cm), Faraday rotation is negligible, so that the measured polarization directly traces the magnetic field in the emitting region. At longer wavelengths, Faraday rotation measurements give additional information on density and magnetic field structure along the entire line of sight. Furthermore, depolarization processes define a distance beyond which polarized radiation is significantly depolarized, which depends on wavelength. So high-frequency measurements probe the total line of sight through the Galaxy, whereas low-frequency polarization observations only trace the nearby part of the ISM.

Specific intriguing small-scale structures and discrete objects have been studied in diffuse polarization observations (Gray et al. 1998, Haverkorn et al. 2000, Uyaniker and Landecker 2002), but the spatial structures have also been analyzed statistically. Simonetti et al. (1984) and Simonetti and Cordes (1986) have studied the structure in Galactic rotation measure $RM$ by comparing $RM$s of polarized extragalactic sources, and of separate components of the same source. The electron density seems to exhibit a power law density structure function (and therefore a power law angular spectrum) (see also Armstrong et al. 1995, Minter and Spangler 1996). Recently, statistical analyses of the diffuse Galactic polarized foreground have been pursued in the form of angular power spectrum studies (Tucci et al. 2000 and 2002, Baccigalupi et al. 2001, Giardino et al. 2002, Bruscoli et al. 2002), with the objective of estimating the importance of the Galactic ISM as a foreground contaminator for CMBR polarization observations (e.g. Seljak 1997, Prunet et al. 2000).

In this paper, we study the statistical properties of the warm ISM and Galactic magnetic field by means of power spectra of Stokes parameters $Q$ and $U$, polarized intensity $P$ and rotation measure $RM$. We also derive the structure function of $RM$ to allow a comparison with earlier studies of $RM$ structure functions from polarized extragalactic sources. Furthermore, by careful selection of reliable $RM$ determinations (i.e. those with low $\chi^2$ of the linear fit to $\phi(\lambda^2)$) in the calculation of structure functions, we can obtain an estimate of how much the structure functions (and power spectra) are influenced by low-quality $RM$s.

We use data from three regions, all at positive Galactic latitudes, in which we observed the diffuse polarized emission at frequencies around 3.5 GHz with the Westerbork Synthesis Radio Telescope (WSRT). For the first two regions, we have multi-frequency observations at a resolution of 5.0′×5.0′ cosec δ. The first region is centered on $(l, b) = (161^\circ, 16^\circ)$ and has a size of about 9′×7′ (Chapter 4), and the second region is centered on $(l, b) = (137^\circ, 7^\circ)$ and is about 7′×7′ in size (Chapter 5). The third region is a part of the Westerbork Northern Sky Survey (WENSS, Rengelink et
al. 1997), a high-resolution radio survey at 327 MHz of the northern hemisphere. For those parts of the WENSS survey that were observed at night, polarization data are usable (Schnitzeler et al., in prep., and Chapter 7). Here we discuss polarization data from the region with $140^\circ \leq l \leq 170^\circ$ and $0^\circ \leq b \leq 30^\circ$.

In Section 8.2 we describe the three sets of data that we analyze in this paper. In Section 8.3 angular power spectrum analysis is introduced, power spectra of the data are presented and discussed and literature data on the angular power spectra is briefly summarized. Section 8.4 gives structure functions for $RM$ in the two multi-frequency measurements. In Section 8.5, the results are discussed, and, finally, in Section 8.6 some conclusions are stated.

8.2 The observations

8.2.1 Multi-frequency WSRT observations

We carried out low frequency radio polarimetry with the Westerbork Synthesis Radio Telescope (WSRT), in two regions of the sky at positive latitudes, in the constellations of Auriga and Horologium. Data were obtained simultaneously at 8 frequencies around 350 MHz with a bandwidth of 5 MHz. Due to radio interference and hardware problems, only data in 5 frequency bands could be used, viz. those centered at the frequencies
341, 349, 355, 360, and 375 MHz. The multi-frequency data allow the study of the frequency dependence of the polarization structure, and the determination of rotation measure \( RM \). We use the technique of mosaicking (i.e. the telescopes cycle through a number of adjacent fields on the sky during a 12hr observation) to obtain a field of view that is larger than the primary beam. Mosaicking also suppresses instrumental polarization to below 1\( \% \) (see Chapter 2). Maps of the Stokes parameters \( I, Q, \) and \( U \) were derived from the observed visibilities. \( RM \)s were computed straightforwardly from the linear relation between polarization angle \( \phi \) and \( \chi^2 \). Because the observed \( RM \) values are small (\( |RM| \leq 10 \text{ rad m}^{-2} \)), there is no 180\( ^\circ \) ambiguity in \( \phi \), and \( RM \)s can be computed with \( |\phi(\lambda_i) - \phi(\lambda_j)| < 90^\circ \) for adjacent wavelengths \( \lambda_i \) and \( \lambda_j \). We define a determination of \( RM \) in a particular beam as reliable if (1) the reduced \( \chi^2 \) of the linear \( \phi(\chi^2) \) - relation is less than 2, and (2) the polarized intensity averaged over wavelength \( \langle P \rangle \) is larger than 20 mJy/beam (i.e. \( \sim 4\sigma \)).

The maximum baseline of the observations was 2700m, yielding a resolution of 1', but smoothing of the Stokes \( Q \) and \( U \) data (using a Gaussian taper in the \((u, v)\)-plane) was applied to obtain a better signal-to-noise. The taper has a value of 0.25 at a baseline value around 300m, where the exact values were chosen so that the beam size is identical at all 5 frequencies, viz. 5.0'\( \times \)5.0' cosec \( \delta \).

The first region, in the constellation Auriga, is centered on \((l, b) = (161^\circ, 16^\circ)\) and is about 9'\( \times \)7' in size. The left panel in Fig. 8.1 shows the polarized intensity \( P \) at 349 MHz in the Auriga region at 5.0' resolution. We do not show the original 1' resolution map, because it is noise-dominated. The maximum polarized brightness temperature in the map is \( T_{b, pol} \approx 13 \text{ K} \), and the noise is about 0.45 K. Rotation measures in the Auriga region are shown in the right panel of Fig. 8.1 as circles superimposed on the grey scale map of \( P \) at 349 MHz. The diameter of each circle indicates the value of the \( RM \) at that position, where filled circles denote positive \( RM \)s, and open circles negative \( RM \)s. We show only the \( RM \)s that we judge to have been reliably determined, and only one in four independent beams.

The region around \((l, b) = (137^\circ, 7^\circ)\) in the constellation Horologium was observed in the same way as the Auriga region, at the same five frequencies, and with the same taper applied. Fig. 8.2 shows the polarized intensity \( P \) at 349 MHz in the left panel, and \( RM \)s in the form of overlaid circles in the right panel. \( RM \)s are coded in the same way as in Fig. 8.1. The maximum polarized brightness temperature is \( T_{b, pol} \approx 17 \text{ K} \), and the noise is a little higher than that in the Auriga field, about 0.65 K.

In both fields there is no small-scale structure visible in the map of total intensity Stokes \( I \), despite the ubiquitous structure on arcminute and degree scales in \( P \). Because the large-scale (\( \geq 1^\circ \)) component of \( I \) cannot be measured with the WSRT due to missing short spacings, we used the Haslam et al. (1981, 1982) radio survey at 408 MHz to estimate the total intensity \( I \) in the Auriga region as \( \sim 34 \text{ K} \) and in the Horologium region as \( \sim 47 \text{ K} \). The polarized intensity \( P \) shows structure on scales from arcminutes to degrees of up to \( \sim 10 - 15 \text{ K} \). Due to the lack of corresponding structure in \( I \), the structure in polarized intensity cannot be entirely due to small-scale synchrotron emission but must be due to other, instrumental and depolarization, mechanisms. For details on observations and analysis, see Chapter 4 for the Auriga region, and Chapter 5 for Horologium.

In a medium that emits synchrotron radiation and simultaneously causes Faraday
rotation, the polarized emission is depolarized by so-called depth depolarization, due to the vector averaging of contributions from different parts of the line of sight. This, together with beam depolarization (due to angle structure within one synthesized beam), produces structure in $P$. Furthermore, the insensitivity of the interferometer to large-scale structure can cause additional structure in $P$. However, we have shown in Chapter 3 that this effect cannot be very important in these observations.

8.2.2 Polarization data from the WENSS survey

The Westerbork Northern Sky Survey (WENSS, Rengelink et al. 1997) is a low-frequency radio survey that covers the whole sky north of $\delta = 30^\circ$ at 325 MHz to a limiting flux density of approximately 18 mJy ($5\sigma$) and with a resolution of $54'' \times 54'' \csc \delta$.

Polarization data taken during the day are greatly affected by solar radiation, which is detected in sidelobes. Furthermore, ionospheric Faraday rotation rapidly changes during sunrise and sunset, causing a reduction of the apparent polarized intensity, and adding much noise. However, in mosaics observed (almost) entirely during night time, the polarization data is of good quality. As a result, we could make a “supermosaic” of a large region of $\sim 30^\circ \times 35^\circ$, which we will refer to as the WENSS polarization region. Details on the observations and correction for ionospheric Faraday rotation are given in Schnitzeler et al. (in prep.) and in Chapter 7.

Fig. 8.3 shows polarized intensity $P$ in the WENSS polarization region, resampled in Galactic longitude and latitude. For this analysis, a Gaussian taper with value 0.25 at a baseline of 500m was applied, which yields a resolution of $\sim 2.5'$. In the figure, $P$ saturates at 35 mJy/beam, which coincides with a polarized brightness temperature of $\sim 17$ K to $\sim 25$ K, depending on declination. The average $P \approx 2.6$ mJy/beam. Note that these observations are taken in a single 5 MHz wide frequency band, so rotation measure data is not available.
Figure 8.3: Grey scale representation of the polarized intensity of the polarization part of the WENSS survey at 325 MHz, where white denotes a maximum intensity of 35 mJy/beam ($T_{b,\text{tot}} \approx$ 17 K to 25 K, depending on declination). The data was smoothed with a 500m taper, and has a resolution of $\approx$ 2.5$'$.

8.3 Angular power spectrum analysis

8.3.1 Determination of the multipole spectral index

To quantify the structure in the polarization maps, we calculate angular power spectra $P_S(\ell)$ as a function of multipole $\ell$. A multipole $\ell$ is a measure of angular scales equivalent to wave number, and is defined as $\ell \approx 180^\circ/\theta$, where $\theta$ is the angular scale in degrees. The angular power spectrum $P_S$ of a radiation field $X$ is the square of the Fourier transform of $X$: $P_S(X, \ell) = |\mathcal{F}(X, \ell)|^2$, where $\mathcal{F}$ denotes the Fourier transform. The observable $X$ can be either Stokes $Q$, Stokes $U$, polarized intensity $P$ or rotation measure $RM$. The power spectra were computed in two dimensions, and averaged over azimuth in radial bins. The multipole spectral index $\alpha$, defined as $P_S(X, \ell) \propto \ell^{-\alpha}$, is calculated from a log-log fit to the power spectrum. In the tapered data, multipoles with higher values of $\ell$ are affected by the tapering, and in the untapered data higher multipoles are dominated by noise. Multipoles with $\ell \leq 200$ correspond to angular scales $\theta \geq 1^\circ$, to which the WSRT is not sensitive.

The visibilities from which a map is made are $V_{\text{map}}(u,v) = V_{\text{obs}}(u,v) T(u,v)$, where $V_{\text{obs}}$ are the observed visibilities, and $T(u,v)$ is the taper function. The calculated
intensity of tapered data is $\mathcal{F}(V_{\text{map}}) = \mathcal{F}(V_{\text{abs}}) * \mathcal{F}(T)$, where $\mathcal{F}$ is a Fourier transform and the asterisk denotes convolution. The power spectrum of the Stokes parameter $X$, $PS(X)$, is then:

$$PS_X(\ell) = |\mathcal{F}(X)|^2 = |\mathcal{F}(X_{\text{abs}})|^2 T^2 = PS_{X,\text{abs}}(\ell) T^2$$  \hspace{1cm} (8.1)

where $X$ is Stokes $I$, $Q$ or $U$. Although polarized intensity $P$ is derived from $Q$ and $U$ and thus not directly observed, correction for the taper in the same way as for $Q$ and $U$ power spectra is a good approximation. As an illustration, Fig. 8.4 shows the power spectrum of $P$ of the tapered data in the Auriga region at 341 MHz (solid line). The dotted line is the same spectrum, but corrected for the tapering according to Eq. (8.1). The power law behavior extends to $\log(\ell) \approx 3.6$.

### 8.3.2 Power spectra from the multi-frequency WSRT studies

In Fig. 8.5, we show the power spectra of $P$ in the Auriga and Horologium regions at 5 frequencies, both for the tapered (upper line) and un tapered data (lower line) in the same plot. The amplitudes of the power spectra of the tapered data are lower than those of the un tapered data only because the intensities are expressed in mJy/beam. Because the beam widths are different for the two datasets, this gives a difference in the magnitude of $P$ in tapered and un tapered data. The power spectra of $RM$ in the Auriga and Horologium field are given in Fig 8.6. Only tapered data give reliable enough $RM$ determinations to produce power spectra for them.

Figs. 8.7 and 8.8 show power spectra for the Stokes parameters $Q$ and $U$ in the Auriga and Horologium region respectively, again for tapered and un tapered data. The corresponding multipole spectral indices $\alpha$, derived for a multipole range of $400 < \ell < 1500$, are given in Table 8.1. At small scales (large $\ell$), the power spectra of the un tapered data flatten out due to the noise in the maps, while the low-resolution data steepen due to the tapering, as illustrated in Fig. 8.4. At large scales, the $Q$ and $U$ power spectra of the tapered data show a decrease, probably due to the lack of large-scale structure (see Section 8.2.1), which is not visible in $P$. Because $P$ is constructed from $\sqrt{Q^2 + U^2}$, the average $P$ is always positive, and the power spectrum of $P$ always shows large-scale structure. Therefore, the decrease in $Q$ and $U$ power spectra at low
Figure 8.5: Power spectra of polarized intensity $P$ for 5 frequency bands in the Auriga region (top) and Horologium region (bottom). In each plot, the upper line of symbols denotes the tapered data, the lower line the untapered data, and the solid lines are linear fits to the spectra. In the untapered data, the spectrum is flattened at high $\ell$ due to noise.

Figure 8.6: Power spectra of $RM$ in the Auriga region (left) and Horologium region (right).
Figure 8.7: Power spectra of Stokes $Q$ (top) and Stokes $U$ (bottom) for 5 frequencies in the Auriga region. Notation as in Fig. 8.5.

ℓ is not visible in $P$. However, if this is the correct explanation for the decrease in $Q$ and $U$ power spectra at low ℓ, the consistency of the slope of the power spectrum of $P$ below ℓ ~ 300 must be fortuitous.

The power spectra of $Q$ and $U$ are steeper and have a larger amplitude than the power spectra of $P$. This could be caused by the presence of a Faraday screen in front of the emitting region. A Faraday screen will rotate the polarization angle, and so induce extra structure in $Q$ and $U$, while leaving $P$ unaltered. This results in a higher amplitude of the power spectrum. As the Faraday screen consists of foreground material, its angular size is large, steepening the spectrum. This effect was also noticed by Tucci et al. (2002).

The logarithmic slope of the power spectrum of polarized intensity is $\alpha_P \approx 2.1 - 2.3$ (Table 8.1), which is slightly higher than most earlier estimates from the literature (although these are taken at higher frequencies, see Section 8.3.4). However, note that the Auriga and Horologium regions were selected for their conspicuous structure in $P$, so we expect these regions to show more structure on large (degree) scales than the “average” ISM, and thus exhibit a steeper spectrum. The power spectra in the Auriga region show somewhat steeper slopes in $Q$ and $U$ than in Horologium, indicating that
the Horologium region probably contains more small-scale structure in the Faraday screen than the Auriga region.

The power spectra of $RM$ in Fig. 8.5 are shallower than the $Q$, $U$ or $P$ power spectra. In fact, we do not expect a direct correspondence between the multipole spectral indices of $RM$ and $P$ (or $Q$, $U$), as the former describes very directly the electron content and magnetic field in the ISM (integrated over the line of sight), whereas in the latter case the polarized radiation is modulated by Faraday rotation and depolarization.

### 8.3.3 Power spectra from the WENSS polarization region

In the WENSS polarization region, power spectra were evaluated for subfields, to be able to study possible dependences of the multipole spectral index on Galactic longitude and/or latitude. The 11 subfields are shown in Fig 8.9, superimposed on grey scale maps of $P$. The power spectra of polarized intensity $P$ are shown in Fig. 8.10, where the subfields are arranged as in Fig 8.9. The power spectra in subfields 9, 10 and 11, at high Galactic latitude $b$, have a lower amplitude than the power spectra at lower $b$, which is consistent with the decreasing amount of $P$ at higher $b$, visible in Fig. 8.3.
Table 8.1: Multipole spectral indices $\alpha$ for observed polarized intensity $P$, Stokes $Q$ and Stokes $U$, for 5 frequencies and their average over frequency, and $\alpha_{RM}$, in the Auriga and Horologium regions. Values for $P$, $Q$ and $U$ are given for tapered data, denoted by a subscript ‘t’, and untapered data. The multipole ranges used to derive $\alpha$ were 400 < $\ell$ < 1900. Only for $\alpha_P$ of the untapered data, the range was smaller because of flattening of the spectrum at higher $\ell$.

<table>
<thead>
<tr>
<th>$\ell$</th>
<th>341 MHz</th>
<th>349 MHz</th>
<th>355 MHz</th>
<th>360 MHz</th>
<th>375 MHz</th>
<th>mean</th>
</tr>
</thead>
<tbody>
<tr>
<td>$\alpha_{P,t}$</td>
<td>2.37 ± 0.19</td>
<td>2.59 ± 0.20</td>
<td>2.38 ± 0.19</td>
<td>2.35 ± 0.17</td>
<td>1.88 ± 0.18</td>
<td>2.32 ± 0.08</td>
</tr>
<tr>
<td>$\alpha_P$</td>
<td>2.06 ± 0.54</td>
<td>2.23 ± 0.46</td>
<td>2.49 ± 0.60</td>
<td>2.00 ± 0.47</td>
<td>2.20 ± 0.59</td>
<td>2.20 ± 0.24</td>
</tr>
<tr>
<td>$\alpha_{Q,t}$</td>
<td>3.55 ± 0.20</td>
<td>3.72 ± 0.21</td>
<td>3.59 ± 0.20</td>
<td>3.43 ± 0.19</td>
<td>2.50 ± 0.20</td>
<td>3.36 ± 0.09</td>
</tr>
<tr>
<td>$\alpha_Q$</td>
<td>3.13 ± 0.24</td>
<td>3.16 ± 0.25</td>
<td>3.16 ± 0.24</td>
<td>3.00 ± 0.24</td>
<td>3.12 ± 0.25</td>
<td>3.12 ± 0.11</td>
</tr>
<tr>
<td>$\alpha_{U,t}$</td>
<td>3.60 ± 0.19</td>
<td>3.80 ± 0.21</td>
<td>3.79 ± 0.21</td>
<td>3.69 ± 0.21</td>
<td>3.49 ± 0.20</td>
<td>3.71 ± 0.09</td>
</tr>
<tr>
<td>$\alpha_U$</td>
<td>3.25 ± 0.24</td>
<td>3.20 ± 0.25</td>
<td>3.23 ± 0.24</td>
<td>3.13 ± 0.24</td>
<td>3.18 ± 0.53</td>
<td>3.20 ± 0.11</td>
</tr>
<tr>
<td>$\alpha_{RM,t}$</td>
<td></td>
<td></td>
<td></td>
<td></td>
<td></td>
<td>0.99 ± 0.08</td>
</tr>
</tbody>
</table>

Horologium

<table>
<thead>
<tr>
<th>$\ell$</th>
<th>341 MHz</th>
<th>349 MHz</th>
<th>355 MHz</th>
<th>360 MHz</th>
<th>375 MHz</th>
<th>mean</th>
</tr>
</thead>
<tbody>
<tr>
<td>$\alpha_{P,t}$</td>
<td>2.11 ± 0.19</td>
<td>2.18 ± 0.19</td>
<td>2.11 ± 0.18</td>
<td>1.98 ± 0.19</td>
<td>1.95 ± 0.19</td>
<td>2.07 ± 0.08</td>
</tr>
<tr>
<td>$\alpha_P$</td>
<td>2.12 ± 0.63</td>
<td>2.36 ± 0.55</td>
<td>2.26 ± 0.66</td>
<td>2.47 ± 0.68</td>
<td>2.49 ± 0.91</td>
<td>2.34 ± 0.31</td>
</tr>
<tr>
<td>$\alpha_{Q,t}$</td>
<td>2.58 ± 0.22</td>
<td>3.17 ± 0.19</td>
<td>2.96 ± 0.19</td>
<td>2.90 ± 0.20</td>
<td>2.78 ± 0.20</td>
<td>2.89 ± 0.09</td>
</tr>
<tr>
<td>$\alpha_Q$</td>
<td>2.15 ± 0.26</td>
<td>2.92 ± 0.22</td>
<td>2.57 ± 0.23</td>
<td>2.60 ± 0.23</td>
<td>2.32 ± 0.25</td>
<td>2.51 ± 0.11</td>
</tr>
<tr>
<td>$\alpha_{U,t}$</td>
<td>2.56 ± 0.22</td>
<td>3.14 ± 0.20</td>
<td>3.04 ± 0.19</td>
<td>3.22 ± 0.20</td>
<td>3.24 ± 0.17</td>
<td>3.04 ± 0.09</td>
</tr>
<tr>
<td>$\alpha_U$</td>
<td>2.07 ± 0.26</td>
<td>3.00 ± 0.22</td>
<td>2.69 ± 0.23</td>
<td>2.85 ± 0.23</td>
<td>2.65 ± 0.25</td>
<td>2.65 ± 0.11</td>
</tr>
<tr>
<td>$\alpha_{RM,t}$</td>
<td></td>
<td></td>
<td></td>
<td></td>
<td></td>
<td>0.94 ± 0.10</td>
</tr>
</tbody>
</table>

Table 8.2: Multipole spectral indices $\alpha$ for observed polarized intensity $P$, Stokes $Q$ and Stokes $U$ in 11 subfields in the WENSS polarization region at 325 MHz.

<table>
<thead>
<tr>
<th>WENSS</th>
<th>($l$, $b$)</th>
<th>($\ell$, $\ell'$)</th>
<th>$\alpha_P$</th>
<th>$\alpha_Q$</th>
<th>$\alpha_U$</th>
</tr>
</thead>
<tbody>
<tr>
<td>field 1</td>
<td>(159, 04)</td>
<td></td>
<td>1.67 ± 0.08</td>
<td>2.55 ± 0.09</td>
<td>2.43 ± 0.10</td>
</tr>
<tr>
<td>field 2</td>
<td>(165, 11)</td>
<td></td>
<td>1.48 ± 0.08</td>
<td>2.10 ± 0.08</td>
<td>2.06 ± 0.08</td>
</tr>
<tr>
<td>field 3</td>
<td>(158, 11)</td>
<td></td>
<td>1.84 ± 0.09</td>
<td>2.50 ± 0.09</td>
<td>2.20 ± 0.09</td>
</tr>
<tr>
<td>field 4</td>
<td>(151, 11)</td>
<td></td>
<td>1.70 ± 0.08</td>
<td>2.09 ± 0.09</td>
<td>2.02 ± 0.09</td>
</tr>
<tr>
<td>field 5</td>
<td>(144, 11)</td>
<td></td>
<td>1.63 ± 0.08</td>
<td>2.28 ± 0.10</td>
<td>2.31 ± 0.10</td>
</tr>
<tr>
<td>field 6</td>
<td>(143, 18)</td>
<td></td>
<td>1.24 ± 0.09</td>
<td>1.96 ± 0.10</td>
<td>2.04 ± 0.10</td>
</tr>
<tr>
<td>field 7</td>
<td>(150, 18)</td>
<td></td>
<td>0.99 ± 0.08</td>
<td>1.44 ± 0.08</td>
<td>1.47 ± 0.09</td>
</tr>
<tr>
<td>field 8</td>
<td>(157, 18)</td>
<td></td>
<td>1.67 ± 0.09</td>
<td>2.33 ± 0.09</td>
<td>2.33 ± 0.09</td>
</tr>
<tr>
<td>field 9</td>
<td>(157, 25)</td>
<td></td>
<td>0.73 ± 0.08</td>
<td>1.24 ± 0.08</td>
<td>1.18 ± 0.09</td>
</tr>
<tr>
<td>field 10</td>
<td>(150, 25)</td>
<td></td>
<td>0.90 ± 0.08</td>
<td>1.18 ± 0.08</td>
<td>1.44 ± 0.08</td>
</tr>
<tr>
<td>field 11</td>
<td>(144, 25)</td>
<td></td>
<td>0.68 ± 0.08</td>
<td>0.77 ± 0.08</td>
<td>0.63 ± 0.08</td>
</tr>
</tbody>
</table>
The multipole spectral indices of the power spectra of $P$, $Q$ and $U$ are given in Table 8.2, and the dependence of $\alpha_P$ on Galactic longitude and latitude is shown in Fig. 8.11. The left plot shows the variation of spectral index with latitude. The observed decrease of spectral index with increasing latitude (i.e. power spectra become flatter with increasing latitude) indicates that there is a decrease in the amount of large-scale structure with increasing latitude, also apparent in Fig. 8.10. The right plot in the figure shows the dependence of spectral index on longitude after rescaling of the data to a standard latitude of $b = 15^\circ$. The fitted slopes in Fig. 8.11, and slopes determined for $\alpha_Q$ and $\alpha_U$ in a similar way, are:

\[
\begin{align*}
\frac{\partial \alpha_P}{\partial b} &= -0.051 \pm 0.004 \\
\frac{\partial \alpha_Q}{\partial b} &= -0.073 \pm 0.004 \\
\frac{\partial \alpha_U}{\partial b} &= -0.066 \pm 0.004 \\
\frac{\partial \alpha_P}{\partial l} &= -0.001 \pm 0.003 \\
\frac{\partial \alpha_Q}{\partial l} &= -0.008 \pm 0.004 \\
\frac{\partial \alpha_U}{\partial l} &= -0.005 \pm 0.004
\end{align*}
\]

In summary, the spectral index $\alpha$ in $P$, $Q$ and $U$ decreases with increasing Galactic latitude, and is consistent with no dependence of $\alpha$ on Galactic longitude. Although the small errors in the derived slopes suggest a good determination of the slope, the large spread of the data points in Fig. 8.11 indicates that a linear gradient is not the perfect model to describe the data.

The decrease of $\alpha$ with latitude indicates that there is more structure in polarization on larger scales at lower latitude. Thus, although small-scale structure in polarization is seen up to very high latitudes, even at frequencies as low as 350 MHz (Katgert and de Bruyn 1999), the amount of structure and its spectral index decrease with latitude,
at least in this region of the sky. The region studied here may be special, as it is situated at the edge of a region of high polarization (the “fan region”, see e.g. Brouw and Spolestra 1976), which is thought to have little structure in magnetic field and/or electron density compared to its surroundings. This high-polarization region extends to \( b \approx 20^\circ \), and its edge could be responsible for the decrease in structure on scales of \( \sim 1^\circ \).

The spectral indices in the Auriga \((b = 16^\circ)\) and Horologium \((b = 7^\circ)\) regions are higher than implied by Fig. 8.11. This might be due to the way in which the Auriga and Horologium regions were selected, viz. because of their remarkable structure on degree scales.

### 8.3.4 Existing literature of power spectra from diffuse polarization

Much work has been done on the determination of power spectra of the diffuse Galactic synchrotron background, because the Galactic synchrotron radiation is a foreground contaminant in Cosmic Microwave Background Radiation (CMBR) polarization measurements at high frequencies \( \nu \approx 30 - 100 \text{ GHz} \). Power spectra of the diffuse polarized synchrotron background intensity have been determined from several radio surveys at frequencies from 408 MHz to 2.7 GHz, in many parts of the sky (Tucci et al. 2000 and
Figure 8.11: Dependence of multipole spectral index $\alpha_P$ on Galactic latitude $b$ (left) and longitude $l$ (right), for 11 subfields in the polarization part of the WENSS survey. The spectral indices in the right plot are normalized to the latitude gradient shown in the left plot.

These power spectra studies are based on the following surveys of polarized radiation:

- **Dwingeloo 25m-dish survey** (Brouw and Spoelstra 1976), of the region $120^\circ < l < 180^\circ$ and $b > -10^\circ$ (although undersampled). This is a multi-frequency survey at 408 MHz, 465 MHz, 610 MHz, 820 MHz and 1411 MHz, with increasing angular resolution of $2.3^\circ$ to $0.5^\circ$.

- **Parkes 2.4 GHz Galactic plane survey** (Duncan et al. 1997), of the region $238^\circ < l < 5^\circ$ and with $|b| < 5^\circ$, at some positions a few degrees higher, at a resolution of $10.4^\circ$.

- **Effelsberg 2.695 GHz Galactic plane survey** (Duncan et al. 1999), of the region $5^\circ < l < 74^\circ$, and $|b| < 5^\circ$, at a resolution of $4.3^\circ$.

- **Effelsberg 1.4 GHz intermediate latitude survey** (Uyaniker et al. 1999), which consists of 4 regions within $45^\circ < l < 210^\circ$ and $-15^\circ < b < 20^\circ$ with an angular resolution of $10.4^\circ$.

- **Australia Telescope Compact Array (ATCA) 1.4 GHz survey** (Gaensler et al. 2001). This is a test region for the Southern Galactic Plane Survey (SGPS, McChure-Griffiths et al. 2001) at $325.5^\circ < l < 332.5^\circ$, $-0.5^\circ < b < 3.5^\circ$, with a resolution of about $1.0^\circ$.

Power spectra of total intensity $I$ and polarized intensity $P$ were derived in these surveys for multipoles over a range of $\ell \approx 10$ to $6000$.

Fig. 8.12 shows the variation of $\alpha_P$ with Galactic longitude, latitude and frequency, using the available data as detailed in the Appendix. In the left plots, the lines show ranges in longitude (top) and latitude (bottom) over which $\alpha_P$ was computed. Solid lines give high multipole numbers ($100 < \ell < 6000$), dashed-dotted lines denote an intermediate multipole range ($30 < \ell < 200$) and the dotted lines give small multipoles ($10 < \ell < 80$). The WSRT data from the Auriga, Horologium en WENSS fields, discussed here, are given in asterisks. In the right plot, $\alpha_P$ against frequency is displayed.
Figure 8.12: Estimates of multipole spectral indices $\alpha_p(\ell)$ from our observations and from the literature, as a function of Galactic longitude (top), Galactic latitude (bottom), and frequency (right). In the left-hand plots, solid lines denote high multipole numbers ($100 < \ell < 6000$), dashed-dotted lines is an intermediate multipole range ($30 < \ell < 200$) and the dotted lines give small multipoles ($10 < \ell < 80$). In the right-hand (frequency) plot, the dotted lines connect observations of the same region made at different frequencies. The WSRT observations discussed in this paper are shown by asterisks.

The connected points are from the same area observed at different frequencies from 408 MHz to 1411 MHz (in the Brouw and Spoelstra paper).

The first conclusion from Fig. 8.12 is that multipole spectral indices vary over the sky from $\alpha \approx 1$ to 3, without showing a clear correlation with Galactic longitude, while only the WENSS subfields show a dependence of spectral index on latitude. However, all surveys have been made at different resolutions and frequencies, and the regions used to compute power spectra are of different sizes. This can explain why a possible dependence of $\alpha$ on latitude was not clearly seen in the other studies. The large variation in slopes of angular power spectra in $P$ indicates that interpretation of the slope is not straightforward, possibly due to large influence of depolarization mechanisms. Care must therefore be taken in extrapolating the results to higher frequencies.

Furthermore, spectral indices show an increase with frequency from 408 MHz to 1.4 GHz. This means that the power spectra become steeper, so the relative amount of small-scale structure decreases. This could be due to the large Faraday rotation at low frequencies. Typical $RM$s of 5 rad m$^{-2}$ are present in the Brouw and Spoelstra survey (Spoelstra 1984), and will rotate polarization angles at 325 MHz by about 250$^\circ$. Variations in $RM$ of a few rad m$^{-2}$ give angle variations of over 90$^\circ$, which would cause beam depolarization if the angle variations occur on scales smaller than the beam (2.3$^\circ$ in this case). Beam depolarization only acts on scales of the synthesized beam, and therefore creates additional structure on small scales in $P$, which flattens the power spectrum. A $\Delta RM$ of 5 rad m$^{-2}$ would cause a variation in polarization angle of about 40$^\circ$ at 820 MHz, and of no more than 10$^\circ$ at 1.4 GHz. So at frequencies above
1.4 GHz, a $\Delta RM$ of 5 rad m$^{-2}$ would cause negligible beam depolarization. In addition, the resolution of the observations generally increases with increasing frequency, which would also cause a decrease in beam depolarization. This might explain why above 1.4 GHz the spectral index does not appear to be correlated with frequency. The fact that spectral indices in the two WSRT regions are much higher than would be expected from this argument can be due to the criteria used to select the two fields.

8.4 Structure functions

The disadvantage of using angular power spectra is that a regular grid of data is required. If the data is very irregularly spaced (e.g. in the case of data from pulsars or extragalactic sources), it is better to use the structure function which, in principle, gives the same information, but can be calculated easily for irregularly spaced data. The structure function $SF$ of a radiation field $X$, as a function of distance lag $d$, is

$$SF_X(d) = \frac{\sum_{i=1}^{N} (X(x_i) - X(x_i + d))^2}{N}$$ (8.2)

where $X(x_i)$ is the value of field $X$ at position $x_i$ and $N$ is the number of data points. If the power spectrum of $RM$ is a power law with spectral index $\alpha$, then the structure function $SF_{RM}$ is (Simonetti et al. 1984)

$$SF_{RM}(d) \propto d^\mu$$ where $\mu = \begin{cases} \frac{\alpha - 2}{2} & \text{for } 2 < \alpha < 4 \\
2 & \text{for } \alpha > 4 \end{cases}$ (8.3)

We determined the structure functions of $RM$ to compare with existing estimates of the structure function of Galactic $RM$ from polarized extragalactic point sources. As the determination of structure functions does not require a regular grid, we can compute the $SF$ including and excluding “bad” data points to examine how the structure functions of $P$ and $RM$ change. This will allow us to estimate the effect of “bad” data in the power spectra in $P$ and $RM$ computed in Section 8.3.

8.4.1 Structure functions of $RM$

Structure functions $SF_{RM}$ in the Auriga and Horologium regions are plotted against distance $d$ in degrees in the log-log plots in Fig. 8.13 (solid line). The minimum distance shown is $d \approx 3\arcmin$ which is the resolution of the $Q$ and $U$ maps. For the evaluation of the SF, only “reliably determined” $RM$ values are used, according to the definition in Section 8.2.1. Although the spectrum is consistent with a flat slope, there is some evidence for a break in the slope at $d = 0.3\arcmin$, primarily in the Horologium field. For larger angular scales, the SF is approximately flat in the Auriga region, and even decreasing in Horologium.

We can estimate the magnitude of the contribution of unreliable determined $RM$s by reevaluating the SF for the complete grid of $RM$ values, instead of only the reliable $RM$s. This estimate is important for a discussion of the power spectra of $RM$, which were evaluated over the complete dataset, including unreliable determined $RM$s. The structure function using the complete dataset is shown in the left panel of Fig. 8.13
Figure 8.13: The solid lines show the \( R \)M structure function \( SF_{RM} \) as a function of distance \( d \) in degrees for the Auriga region (left) and the Horologium region (right), using only reliable \( RM \) values. The dotted line in the left plot gives \( SF_{RM} \) evaluated for the entire grid, including unreliable \( RM \) determinations.

![Auriga and Horologium plots](image)

Figure 8.14: Structure function \( SF_P \) as a function of distance \( d \) in degrees for the Auriga region, in 5 frequencies. The solid line shows \( SF_P \) in which only beams with \( P > 4\sigma \) are included, the dotted line gives \( SF_P \) as computed from all data.

![Auriga plot](image)

as a dotted line. The structure function clearly has a lower amplitude if the unreliable \( RM \) determinations are removed, but the slope of the structure function remains approximately the same.

### 8.4.2 Structure functions of \( P \)

We compute the structure function of polarized intensity, for both the complete grid of beams, and for those beams selected to have high \( P \). In Fig. 8.14 we show structure functions of \( P \) in the Auriga region for 5 frequencies. The average logarithmic slope of the structure functions is \( \sim 0.35 \) in the range \( 0.2^\circ \lesssim d \lesssim 1^\circ \), and the spectrum flattens on larger scales, probably due to the insensitivity of the WSRT to structure on large angular scales. The solid line denotes the estimate based only on beams with \( P > 4\sigma \), and the dotted line gives the estimate based on all the data. The structure functions
based on selected beams and those based on all data are not significantly different on small scales, and start deviating only for $d \gtrsim 0.7$. Although small deviations are created by selecting the best data, the overall slope of the structure function is hardly affected by the selection including only regions of high $P$. The relatively small effect of including bad data gives confidence that the determination of the logarithmic slope of the power spectrum of $RM$ and $P$ is not significantly influenced by noisy data or poorly determined $RM$.

8.5 Discussion

As far as we know, the results presented here are the first observational determinations of the power spectrum of rotation measure. What power spectrum is to be expected is unclear, as the $RM$ is a complex quantity depending on magnetic field structure, electron density and the path length through the ionized ISM. There are indications that the electron density and $RM$ exhibit Kolmogorov turbulence (Minter and Spangler 1996, Armstrong et al. 1993).

An approximately flat structure function corresponds to a structure of equal amplitude on all scales, similar to a noise spectrum. The break in $SF_{RM}$ at $d\theta \approx 0.3^\circ$, if present, corresponds to a change in characteristics of the structure on scales of $\sim 3.9$ pc, assuming a path length of $600$ pc (Chapter 3).

Minter and Spangler (1996) studied structure functions of $RM$ from polarized extragalactic sources and of emission measure (EM) from H$\alpha$ measurements, on the same angular scales as in our observations. They find a break in the slope of the structure function, which can be interpreted as a transition of 3D Kolmogorov turbulence ($\mu = 5/3$, where $\mu$ is the slope of the structure function as defined in Eq. (8.3)) to 2D turbulence ($\mu = 2/3$) in $RM$ on angular scales of $d\theta \approx 0.1^\circ$, the scale of our resolution. They assume a total path length of $2000$ pc, and conclude that the transition occurs at scales of $\sim 3.6$ pc, consistent with the spatial scale at which we estimate the turnover point in the SF of the Auriga and Horologium regions. However, there is a large discrepancy between $\mu$ values found by Minter and Spangler ($\mu = 5/3$ and $2/3$) and by us ($\mu \approx 0$), which could be explained by the fact that they probe the complete line of sight through the medium of many kpc, whereas the $RM$ that we obtain is only produced in the nearest $\sim 600$ pc. Therefore, it is possible that the nature of the turbulence changes from $\mu = 0$ nearby (approximately in the Galactic stellar disk), to 2D and/or Kolmogorov-like turbulence at larger distances.

However, flat structure functions are also found by Simonetti et al. (1984), who have studied structure functions from $RM$s in three regions of the sky. Two of the three regions (around the North Galactic Pole and at $180^\circ < l < 220^\circ$, $10^\circ < b < 50^\circ$) are consistent with a flat structure function ($\mu = 0$). (The third region is located along the Local arm and in the plane, at $70^\circ < l < 110^\circ$ and $-45^\circ < b < 5$, and its structure function of $RM$ has a slope $\mu \approx 1$.) The smallest angular scale they probe in these regions is about $2^\circ$. Clegg et al. (1992) study low-latitude extragalactic sources and also obtain flat structure functions, albeit with a much higher amplitude than that of the high-latitude extragalactic sources.
8.6 Conclusions

The multipole spectral index for polarized intensity $P$ is $\alpha \approx 2.2$ in the Auriga and Horologium regions, and ranges from 0.7 to 1.8 for subfields in the WENSS polarization region. The multipole spectral index decreases with Galactic latitude (i.e., power spectra become flatter towards higher latitudes), but is probably constant with Galactic longitude.

In all regions, the power spectra of Stokes $Q$ and $U$ are steeper than the spectra of $P$. This is most likely due to the presence of a Faraday screen, which creates additional structure in $Q$ and $U$, but not in $P$. As the Faraday screen is located in front of the polarized emission, the structure induced by the screen will be on larger angular scales than that of the emission, which steepens the $Q$ and $U$ spectra.

The derived power spectra of $P$ agree with earlier estimates, although all estimates show a large range of $0.6 \leq \alpha_P \leq 3$, possibly due to a large influence of depolarization mechanisms. This makes interpretation of the power spectra uncertain.

Structure functions of $RM$ in the Auriga and Horologium fields are consistent with flat spectra (i.e., a logarithmic slope of the structure function $\mu = 0$), but may show a break close to $0.3'$, which is at the same spatial scales as a break in the structure function in the $RMs$ of extragalactic sources (Minter and Spangler 1996). The flat spectrum indicates a noise-like spectrum with equal amounts of structure on all scales.

The derived structure functions support the estimates of the power spectra, because they are based only on high-quality data, i.e. reliably determined $RM$, and $P$ with high S/N. Structure functions using all data do not show a significant difference with the structure functions of the selected data. This gives confidence that the power spectra determinations are not significantly affected by bad data points.

Acknowledgements

We thank B. J. Rickett for helpful discussions and suggestions, and F. Heitsch for help in constructing and analyzing power spectra. The Westerbork Synthesis Radio Telescope is operated by the Netherlands Foundation for Research in Astronomy (ASTRON) with financial support from the Netherlands Organization for Scientific Research (NWO). Computations presented here were performed on the SGI Origin 2000 machine of the Rechenzentrum Garching of the Max-Planck-Gesellschaft. MH acknowledges the support from NWO grant 614-21-006.

References

Ferrière K. M., 2001, RVMP 73, 1031
Landecker T. L., Uyaniker B., & Kothes R., 2001, AAS 199, #58.07
Tucci M., Carretti E., Cecchini S., Fabbri R., Orsini M., & Pierpaoli E., 2000, NewA 5, 181
Appendix

<table>
<thead>
<tr>
<th>survey</th>
<th>$l$</th>
<th>$b$</th>
<th>$\nu$ (MHz)</th>
<th>$\ell$</th>
<th>$\alpha_P$</th>
<th>ref.</th>
</tr>
</thead>
<tbody>
<tr>
<td>Dwingeloo</td>
<td>110 - 160</td>
<td>0 - 20</td>
<td>1411</td>
<td>30 - 100</td>
<td>2.9</td>
<td>(1)</td>
</tr>
<tr>
<td>Brouw &amp; Spoelstra (1976)</td>
<td>5 - 80</td>
<td>50 - 90</td>
<td>&quot;&quot;</td>
<td>&quot;&quot;</td>
<td>3.1</td>
<td></td>
</tr>
<tr>
<td></td>
<td>335 - 360</td>
<td>60 - 90</td>
<td>&quot;&quot;</td>
<td>&quot;&quot;</td>
<td>2.8</td>
<td></td>
</tr>
<tr>
<td></td>
<td>120 - 180</td>
<td>-10 - 20</td>
<td>408</td>
<td>10 - 70</td>
<td>1.3</td>
<td>(2)</td>
</tr>
<tr>
<td></td>
<td>&quot;&quot;</td>
<td>&quot;&quot;</td>
<td>465</td>
<td>&quot;&quot;</td>
<td>1.1</td>
<td></td>
</tr>
<tr>
<td></td>
<td>&quot;&quot;</td>
<td>&quot;&quot;</td>
<td>610</td>
<td>&quot;&quot;</td>
<td>1.5</td>
<td></td>
</tr>
<tr>
<td></td>
<td>&quot;&quot;</td>
<td>&quot;&quot;</td>
<td>820</td>
<td>&quot;&quot;</td>
<td>1.5</td>
<td></td>
</tr>
<tr>
<td></td>
<td>&quot;&quot;</td>
<td>&quot;&quot;</td>
<td>1411</td>
<td>&quot;&quot;</td>
<td>1.8</td>
<td></td>
</tr>
<tr>
<td></td>
<td>45 - 75</td>
<td>26 - 62</td>
<td>408</td>
<td>10 - 70</td>
<td>1.3</td>
<td></td>
</tr>
<tr>
<td></td>
<td>&quot;&quot;</td>
<td>&quot;&quot;</td>
<td>465</td>
<td>&quot;&quot;</td>
<td>1.3</td>
<td></td>
</tr>
<tr>
<td></td>
<td>&quot;&quot;</td>
<td>&quot;&quot;</td>
<td>610</td>
<td>&quot;&quot;</td>
<td>1.6</td>
<td></td>
</tr>
<tr>
<td></td>
<td>&quot;&quot;</td>
<td>&quot;&quot;</td>
<td>820</td>
<td>&quot;&quot;</td>
<td>1.7</td>
<td></td>
</tr>
<tr>
<td></td>
<td>&quot;&quot;</td>
<td>&quot;&quot;</td>
<td>1411</td>
<td>&quot;&quot;</td>
<td>1.7</td>
<td></td>
</tr>
<tr>
<td></td>
<td>15 - 45</td>
<td>60 - 84</td>
<td>408</td>
<td>10 - 70</td>
<td>1.1</td>
<td></td>
</tr>
<tr>
<td></td>
<td>&quot;&quot;</td>
<td>&quot;&quot;</td>
<td>465</td>
<td>&quot;&quot;</td>
<td>0.9</td>
<td></td>
</tr>
<tr>
<td></td>
<td>&quot;&quot;</td>
<td>&quot;&quot;</td>
<td>610</td>
<td>&quot;&quot;</td>
<td>1.2</td>
<td></td>
</tr>
<tr>
<td></td>
<td>&quot;&quot;</td>
<td>&quot;&quot;</td>
<td>820</td>
<td>&quot;&quot;</td>
<td>1.2</td>
<td></td>
</tr>
<tr>
<td></td>
<td>&quot;&quot;</td>
<td>&quot;&quot;</td>
<td>1411</td>
<td>&quot;&quot;</td>
<td>1.9</td>
<td></td>
</tr>
</tbody>
</table>

| Parkes                     | 240 - 250 | -5 - 5 | 2400       | 100 - 800 | 1.86       | (1)  |
| Duncan et al. (1997)       | 250 - 260 | ""    | ""        | ""       | 1.86       |      |
|                            | 260 - 270 | ""    | ""        | ""       | 1.43       |      |
|                            | 270 - 280 | ""    | ""        | ""       | 1.67       |      |
|                            | 280 - 290 | ""    | ""        | ""       | 1.91       |      |
|                            | 290 - 300 | ""    | ""        | ""       | 1.77       |      |
|                            | 300 - 310 | ""    | ""        | ""       | 1.28       |      |
|                            | 310 - 320 | ""    | ""        | ""       | 1.64       |      |
|                            | 320 - 330 | ""    | ""        | ""       | 1.93       |      |
|                            | 330 - 340 | ""    | ""        | ""       | 1.44       |      |
|                            | 340 - 350 | ""    | ""        | ""       | 1.47       |      |
|                            | 350 - 360 | ""    | ""        | ""       | 1.49       |      |
|                            | 240 - 360 | ""    | ""        | ""       | 1.7        | (2)  |
|                            | 326.5 - 331.5 | -1 - 4 | "" | ""       | 1.68        | (3)  |
|                            | 240 - 360 | -5 - 5 | 40 - 250  | 2.4       |            |      |

Table 8.3: Studies of power spectra of diffuse polarized intensity from different surveys. Given are the range in Galactic longitude $l$ and latitude $b$, the frequency of observation $\nu$, the range in multipole $\ell$ for which the power spectrum was computed, the spectral index of $P \alpha_P$ and the reference to the study of the power spectra.
<table>
<thead>
<tr>
<th>survey</th>
<th>( l )</th>
<th>( b )</th>
<th>( \nu ) (MHz)</th>
<th>( \ell )</th>
<th>( \alpha_P )</th>
<th>ref.</th>
</tr>
</thead>
<tbody>
<tr>
<td>Effelsberg</td>
<td>20 - 30</td>
<td>- 5 - 5</td>
<td>2695</td>
<td>100 - 800</td>
<td>1.79</td>
<td>(1)</td>
</tr>
<tr>
<td>Duncan et al.</td>
<td>30 - 40</td>
<td>&quot;</td>
<td>&quot;</td>
<td>&quot;</td>
<td>1.72</td>
<td></td>
</tr>
<tr>
<td>(1999)</td>
<td>40 - 50</td>
<td>&quot;</td>
<td>&quot;</td>
<td>&quot;</td>
<td>1.55</td>
<td></td>
</tr>
<tr>
<td></td>
<td>50 - 60</td>
<td>&quot;</td>
<td>&quot;</td>
<td>&quot;</td>
<td>1.98</td>
<td></td>
</tr>
<tr>
<td></td>
<td>55 - 65</td>
<td>&quot;</td>
<td>&quot;</td>
<td>&quot;</td>
<td>1.93</td>
<td></td>
</tr>
<tr>
<td></td>
<td>5 - 75</td>
<td>&quot;</td>
<td>&quot;</td>
<td>&quot;</td>
<td>1.6</td>
<td>(2)</td>
</tr>
<tr>
<td>Effelsberg</td>
<td>45 - 55</td>
<td>5 - 20</td>
<td>1400</td>
<td>100 - 800</td>
<td>1.5</td>
<td>(2)</td>
</tr>
<tr>
<td>Uyaniker et al.</td>
<td>140 - 150</td>
<td>4 - 10</td>
<td>&quot;</td>
<td>&quot;</td>
<td>2.5</td>
<td></td>
</tr>
<tr>
<td>(1999)</td>
<td>190 - 200</td>
<td>8 - 15</td>
<td>&quot;</td>
<td>&quot;</td>
<td>2.3</td>
<td></td>
</tr>
<tr>
<td>ATCA</td>
<td>327 - 331</td>
<td>- 0.5 - 3.5</td>
<td>1400</td>
<td>600 - 6000</td>
<td>1.68</td>
<td>(3)</td>
</tr>
<tr>
<td>Gaensler et al.</td>
<td>329 - 332</td>
<td>0 - 3</td>
<td>&quot;</td>
<td>&quot;</td>
<td>1.66</td>
<td></td>
</tr>
<tr>
<td>(2001)</td>
<td>326 - 329</td>
<td>&quot;</td>
<td>&quot;</td>
<td>&quot;</td>
<td>1.68</td>
<td></td>
</tr>
<tr>
<td></td>
<td>330.1 - 331.1</td>
<td>0.5 - 1.5</td>
<td>&quot;</td>
<td>&quot;</td>
<td>1.79</td>
<td></td>
</tr>
<tr>
<td></td>
<td>329.2 - 330.2</td>
<td>0.6 - 1.6</td>
<td>&quot;</td>
<td>&quot;</td>
<td>2.24</td>
<td></td>
</tr>
<tr>
<td></td>
<td>327.2 - 328.2</td>
<td>&quot;</td>
<td>&quot;</td>
<td>&quot;</td>
<td>1.93</td>
<td></td>
</tr>
<tr>
<td></td>
<td>328 - 329</td>
<td>2.1 - 3.1</td>
<td>&quot;</td>
<td>&quot;</td>
<td>2.58</td>
<td></td>
</tr>
<tr>
<td></td>
<td>326.5 - 327.5</td>
<td>1.8 - 2.8</td>
<td>&quot;</td>
<td>&quot;</td>
<td>1.88</td>
<td></td>
</tr>
<tr>
<td></td>
<td>329.9 - 330.9</td>
<td>&quot;</td>
<td>&quot;</td>
<td>&quot;</td>
<td>1.68</td>
<td></td>
</tr>
<tr>
<td>WSRT</td>
<td>158 - 165</td>
<td>13 - 20</td>
<td>350</td>
<td>100 - 1000</td>
<td>2.26</td>
<td>Ch. 4</td>
</tr>
<tr>
<td>WSRT</td>
<td>134 - 141</td>
<td>3 - 10</td>
<td>350</td>
<td>100 - 1000</td>
<td>2.20</td>
<td>Ch. 5</td>
</tr>
<tr>
<td>WSRT</td>
<td>156 - 143</td>
<td>0.5 - 7.5</td>
<td>327</td>
<td>100 - 1500</td>
<td>1.67</td>
<td>Ch. 7</td>
</tr>
<tr>
<td></td>
<td>162 - 169</td>
<td>8 - 15</td>
<td>&quot;</td>
<td>&quot;</td>
<td>1.48</td>
<td></td>
</tr>
<tr>
<td></td>
<td>155.5 - 161.5</td>
<td>&quot;</td>
<td>&quot;</td>
<td>&quot;</td>
<td>1.84</td>
<td></td>
</tr>
<tr>
<td></td>
<td>147.5 - 154.5</td>
<td>&quot;</td>
<td>&quot;</td>
<td>&quot;</td>
<td>1.70</td>
<td></td>
</tr>
<tr>
<td></td>
<td>140.5 - 147.5</td>
<td>&quot;</td>
<td>&quot;</td>
<td>&quot;</td>
<td>1.63</td>
<td></td>
</tr>
<tr>
<td></td>
<td>130 - 146</td>
<td>15 - 22</td>
<td>&quot;</td>
<td>&quot;</td>
<td>1.24</td>
<td></td>
</tr>
<tr>
<td></td>
<td>146.5 - 153.5</td>
<td>&quot;</td>
<td>&quot;</td>
<td>&quot;</td>
<td>0.99</td>
<td></td>
</tr>
<tr>
<td></td>
<td>153.5 - 160.5</td>
<td>&quot;</td>
<td>&quot;</td>
<td>&quot;</td>
<td>1.67</td>
<td></td>
</tr>
<tr>
<td></td>
<td>153.3 - 160.5</td>
<td>22 - 29</td>
<td>&quot;</td>
<td>&quot;</td>
<td>0.73</td>
<td></td>
</tr>
<tr>
<td></td>
<td>147 - 154</td>
<td>&quot;</td>
<td>&quot;</td>
<td>&quot;</td>
<td>0.90</td>
<td></td>
</tr>
<tr>
<td></td>
<td>140.5 - 147.5</td>
<td>&quot;</td>
<td>&quot;</td>
<td>&quot;</td>
<td>0.68</td>
<td></td>
</tr>
</tbody>
</table>

(1) Baccigalupi et al. 2001
(2) Bruscoli et al. 2002
(3) Tucci et al. 2002
(4) Giardino et al. 2002

Table 8.3 Continued
MHD simulations of the warm ISM, and comparison with observations: preliminary results

M. Haverkorn and F. Heitsch

Abstract

We present numerical simulations of the warm ionized gaseous component of the Interstellar Medium (ISM), and compare those with radio polarimetric observations. A full-scale 3-dimensional magnetohydrodynamic (MHD) model was constructed which contains a single-component gas and magnetic field. At this stage we only present a low-resolution model, which limits the predictive power of the model. Therefore, we constructed a complementary set of static models, developed as realizations of independent power law distributions of magnetic field density and electron density. To be able to compare the predictions of the models with observations of the polarized diffuse radio background, the model is irradiated with a polarized radio background. The polarization angle is Faraday-rotated while the radiation propagates through the medium. The depolarization process of beam depolarization, i.e. the decrease in polarization due to averaging different polarization vectors across the telescope beam, is simulated. This is done by convolution of the “observables” Stokes $Q$ and $U$ with Gaussian “telescope beams” of different widths. Because the simulated medium acts as a Faraday screen, beam depolarization is the only depolarization mechanism present. The simulations are compared to the observations by generating spatial distributions and power spectra of Stokes $Q$ and $U$, polarized intensity $P$ and rotation measure $RM$. Static models with steep power spectra in magnetic field strength (spectral index $\geq 2.5$) fail to reproduce the observations. The apparent $P$ and $RM$ maps constructed from smoothed $Q$ and $U$ maps in the dynamical model show characteristics similar to the maps in the observations. This indicates the importance of beam depolarization in the creation of one-beam wide “canals” of depolarization. The degree of polarization decreases slowly with increased smoothing in the dynamical models, so structure on scales much smaller than the beam size is needed to obtain the observed degree of depolarization by beam depolarization only. As structure on such small scales does not appear to be dominant in the observations, this indicates that other depolarization mechanisms are important in the observations as well.

161
9.1 Introduction

The Galactic interstellar medium (ISM) is a complex multi-component medium, which can be studied in many ways. Both observational and numerical studies have shed light on the components of and physical processes in the ISM.

Numerical studies suggest a much more complex ISM than the relatively simple three-phase medium in pressure equilibrium that has emerged from early models. Instead, the ISM appears to be a highly turbulent and transient medium. Hydrodynamical and magnetohydrodynamical simulations have elucidated the pressure balance in the ISM, the heating and formation of molecular clouds, etc. (see e.g. reviews by Vázquez-Semadeni 2002 and Mac Low 2002). However, due to the complexity of the interstellar medium, existing numerical simulations only probe a part of the ISM, and/or do not include all the relevant physical processes. A full model of the ISM, including self-gravity, realistic turbulent driving, chemistry, stellar feedback, Galactic rotation, magnetic fields, cosmic rays etc. does not exist as yet. Moreover, direct comparisons between numerical results and astrophysical observations are scarce, and mostly concentrate on molecular clouds (e.g. Falgarone et al. 1994; Padoan et al. 1999; Rosolowsky et al. 1999, MacLow and Ossenkopf 2000, Ossenkopf and MacLow 2002).

We make a first attempt in this chapter to compare simulations of the warm ionized part of the ISM to polarimetric observations of the diffuse radio background. The cosmic rays pervading the Galaxy emit synchrotron radiation, and so provide a partially polarized radio background. All high-resolution observations performed in the last decade (e.g. Wieringa et al. 1993, Duncan et al. 1997, 1999, Gray et al. 1998, Uyaniker et al. 1999, Gaensler et al. 2001, Uyaniker and Landecker 2002, this thesis) show abundant small-scale structure in polarized intensity and polarization angle, often not correlated to structure in total intensity. This led Wieringa et al. (1993) to conclude that the small-scale structure in the polarization angle must be caused by Faraday rotation in a magneto-ionic medium. The Faraday rotation of the polarization angle \( \phi \) is proportional to wavelength squared. The proportionality constant is the rotation measure \( RM = 0.81 \int n_e B_\parallel ds \) where \( n_e \) is the electron density in \( \text{cm}^{-3} \), \( B_\parallel \) the parallel component of the Galactic magnetic field in \( \mu \text{G} \), and \( ds \) the path length through the magneto-ionic medium in parsecs. As Faraday rotation only modifies the polarization angle, it cannot be the cause for the structure in polarized intensity \( P \). Instead, the small-scale structure in \( P \) must be the result of depolarization mechanisms, viz. beam depolarization and depth depolarization (see Chapter 3).

We define beam depolarization as the (partial) canceling of different polarization vectors across the telescope beam. Depth depolarization is defined as the process that cancels different polarization vectors along the line of sight. These definitions are not related to the physical process which cause the variation in polarization vectors. Instead, they give a purely geometrical division of depolarization processes, which is most convenient in this analysis. A more physical description of depolarization is based on the different physical processes causing change of polarization vectors (Burn 1966, Sokoloff et al. 1998, and Chapter 3 for a short summary).

In this chapter, we model the warm ionized ISM (WIM) as a Faraday screen. So the medium contains thermal electrons and magnetic field, and therefore causes Faraday
rotation, but it does not emit synchrotron radiation. The medium is irradiated by a constant polarized background, which results in a constant total intensity \( I \) and constant polarized intensity \( P \), but small-scale structure in Stokes parameters \( Q \) and \( U \) is produced as a result of Faraday rotation in the screen. A background of constant \( I \) is implied by the lack of structure in the interferometric \( I \) image (see Section 9.4).

In the models, structure in polarized intensity is created by depolarization. As the simulated medium does not emit synchrotron radiation, depth depolarization cannot occur. Therefore, all structure in \( P \) in the simulations is caused by beam depolarization, which is modeled by convolution of the “observables” Stokes \( Q \) and \( U \) with a Gaussian function of width \( \sigma \). This simulates structure in polarization angle across the telescope beam on scales of \( 1/\sigma \) times the beam width.

At this stage of the project, the models do not yet allow full quantitative comparison with the observations. Therefore, this paper is meant to give a first impression of the possibilities of these models, rather than to yield quantitative conclusions about the ISM.

### 9.2 The numerical simulations

Ideally, we would like to compare the observations to full-scale 3-dimensional magneto-hydrodynamic (MHD) models of the warm ionized ISM. However, the high resolution and large field of view that are available in the observations pose severe challenges to any numerical simulation. Thus, we constructed two types of models: first, we evolved the full set of time-dependent MHD equations within a bar-like volume, resembling the observational domain. Magnetic field and gas density are coupled self-consistently. However, the model has much lower resolution than the observations. To overcome the resolution problem, we constructed static models, prescribing density and magnetic field structure according to a power law. Although these can reach resolutions comparable to the observations, the density and magnetic field are not coupled, which has consequences for the resulting \( RM \), as explained below.

<table>
<thead>
<tr>
<th>model</th>
<th>( s1 )</th>
<th>( s2 )</th>
<th>( s3 )</th>
<th>( s4 )</th>
<th>( s5 )</th>
<th>( d1 )</th>
</tr>
</thead>
<tbody>
<tr>
<td>type</td>
<td>s</td>
<td>s</td>
<td>s</td>
<td>s</td>
<td>s</td>
<td>s</td>
</tr>
<tr>
<td>resolution</td>
<td>512(^2)</td>
<td>512(^3)</td>
<td>512(^3)</td>
<td>512(^3)</td>
<td>512(^3)</td>
<td>128 \times 128 \times 1024</td>
</tr>
<tr>
<td>( L ) [pc]</td>
<td>600</td>
<td>600</td>
<td>600</td>
<td>600</td>
<td>600</td>
<td>400</td>
</tr>
<tr>
<td>( B ) [( \mu )G]</td>
<td>3.0</td>
<td>1.5</td>
<td>1.5</td>
<td>0.7</td>
<td>0.3</td>
<td>2.0</td>
</tr>
<tr>
<td>( n_e ) [cm(^{-3})]</td>
<td>0.1</td>
<td>0.1</td>
<td>0.1</td>
<td>0.1</td>
<td>0.1</td>
<td>0.1</td>
</tr>
<tr>
<td>( \alpha )</td>
<td>0.0</td>
<td>1.6</td>
<td>2.0</td>
<td>2.6</td>
<td>3.4</td>
<td>*</td>
</tr>
</tbody>
</table>

**Table 9.1:** Model parameters for static (type s) and dynamical (type d) models. Physical length is \( L \), \( B \) gives the rms field strength and \( n_e \) is the mean electron density. The power spectra of magnetic energy density \(( \propto B^2 \) in the static models are defined as \( P_S(k) \propto k^{-2\alpha} \).
9.2.1 Dynamical model

In order to evolve the MHD equations, we used the well-tested 3-dimensional Eulerian finite difference code ZEUS-3D (Stone and Norman 1992a,b, Hawley and Stone 1995). The MHD-equations describe the time evolution of a single-component charge-neutral gas under the influence of thermal pressure and magnetic tension and pressure forces. The magnetic field follows the induction equation with zero resistivity, i.e. under the assumption of perfect flux freezing. Numerically, this assumption breaks down at grid scale, which provides the energy dissipation channel at the small-scale end of the turbulent cascade.

We model the turbulence in the WIM by external forcing. Each time step, we add a fixed grid of perturbations to the three velocity components. The perturbations are chosen from a Gaussian random distribution in Fourier space such that only the largest scales in a shell between wavenumbers \(1 < kL/2\pi < 2\), where \(L\) is the box length, are excited. The amplitudes are normalized each timestep so that the energy input rate is constant, thus mimicking driving on the largest possible scales (Mae Low and Ossenkopf, 2000) and a continuous flow of turbulent kinetic energy from larger scales into the computational domain. After a dynamical time, the full turbulent cascade has developed, and the model is ready for investigation. The resulting power spectrum shows the three ranges typical for a turbulence spectrum: the source region at the largest scales, the inertial range (i.e. the range where a power law applies) for intermediate scales and the dissipative range at the smallest scales. The latter is purely numerical, and the goal of any turbulence simulation is to achieve sufficiently high resolution to identify the inertial range unambiguously.

In an attempt to mimic the observational domain as closely as possible, we constructed models with bar-like domains of resolution \(128 \times 128 \times 1024\) cells, where the line of sight is parallel to the long axis of the bar. As the turbulence driver relies on wave number isotropy, we have eight statistically independent turbulence cells along the line of sight. Furthermore, we cannot introduce structure at scales larger than the lowest wavenumber, a restriction which has to be kept in mind when comparing to observations.

We use the ideal gas law as an isothermal equation of state with temperature \(T = 8000\) K, and a rms Mach number of \(\mathcal{M} = 1\) in the dynamically evolved state. The particle density is \(n = 0.1\) cm\(^{-3}\). The initially uniform magnetic field of strength \(B = 2\) \(\mu\)G is oriented perpendicularly to the line of sight.

The inertial range of the turbulence spectrum is strongly influenced by numerical dissipation, due to the limited resolution. This limits the applicability of the model. Thus we decided to construct a complementary set of static models.

9.2.2 Static models

Static models are much less expensive in terms of computation time, so that we can construct them at higher resolution (\(512^3\) cells so far) and for a range of parameters. We use these models in the first place to determine the quality of \(RM\) determinations. They are certainly not a full-scale numerical counterpart to observations.

We assume an isotropic power law spectrum for the magnetic field energy density
and an independent power law for the electron density, so that magnetic field and electron density are not correlated. To assure that the magnetic field is divergence-free, we actually constructed a cube of isotropic vector potential $A$, realized from a power law with slope $\alpha_B + 1$ and with random phases. The magnetic field $B = \nabla \times A$ yields an isotropic, divergence-free magnetic field according to a power spectrum with slope $\alpha_B$. We constructed models with different slopes of power spectra of the magnetic field, viz. $\alpha_B = 0, -1.6, -2.0, -2.6$ and $-3.4$ (see Table 9.1).

The thermal electron density is constrained by the lack of structure in emission measure obtained from Hα measurements in the radio polarimetry fields (Havercorn et al., in prep.), conducted with the Wisconsin Hα Mapper (WHAM, Hafner et al., in prep., Reynolds et al. 1998). We assumed a power law with slope $\alpha_n$, equal to the power law slope of the magnetic field strength, and the amplitude is set in agreement with the upper limits set by the Hα observations so that $0.97 < n_e < 1.03 \text{ cm}^{-3}$.

Because in the static models the input magnetic field and electron density are unrelated and there is no natural scale, the magnetic field, electron density and path length can be scaled arbitrarily. We assumed a path length of 600 pc, approximately equal to the estimated path length of the observations (Chapter 3), the electron density determined by the constraint of the Hα measurements, and adjusted the strength of the magnetic field so that the output $RM$ is comparable to the $RM$ from the observations, see Table 9.1.

### 9.2.3 Using the modeled medium as a Faraday screen

As mentioned in Section 9.1, the medium is a Faraday screen irradiated by a polarized background of radio emission, which is Faraday-rotated in passage through the medium. Because the radio polarimetric observations discussed in this chapter show a constant total intensity (Section 9.4), we irradiate the models with a 100% polarized background without intrinsic structure.

Faraday rotation causes structure in polarization angle. The polarization angle follows from $\phi = \phi_0 + RM \lambda^2$, where $\phi_0$ is the intrinsic polarization angle of the background radiation, which is assumed to be zero. The $RM$ is computed by integrating the parallel component of the magnetic field multiplied by the electron density over the line of sight. The medium is irradiated by a polarized radio background at five wavelengths $\lambda = 80, 83, 84, 86$ and 88 cm, equal to the wavelengths of the observations.

From the polarization angle $\phi$, the Stokes parameters $Q$ and $U$ are computed according to $Q = P_0 \cos(2\phi)$ and $U = P_0 \sin(2\phi)$, where $P_0$ is the background polarized intensity, scaled to 1. Therefore, the “observables” Stokes $Q$ and $U$ and polarization angle $\phi$ have small-scale structure depending on $RM$, but polarized intensity $P = \sqrt{Q^2 + U^2} = P_0$ is constant.

### 9.2.4 Simulation of beam depolarization by smoothing

The process of beam depolarization in the observations can be simulated by smoothing the Stokes $Q$ and $U$ from the numerical models with a “telescope beam”. The beam is approximated by convolution of the data with a Gaussian function of width $\sigma$. This simulates the situation of an observation with a fixed telescope beam, in situations with
structure across the beam on scales of $1/\sigma$ times the beam width. From the smoothed maps of Stokes $Q$ and $U$, viz. $Q_x$ and $U_x$, polarized intensity $P_z = \sqrt{Q_x^2 + U_x^2}$ and polarization angle $\phi_x = 0.5\arctan(U_x/Q_x)$ can be computed. A linear fit of $\phi_x$ against $\lambda^2$ then yields the $RM$ value as it would have been observed.

9.3 Results from the simulations

To be able to compare the results of the two types of models to the observations, we constructed maps of apparent polarized intensity $P_z$ and apparent rotation measure $RM_x$ after smoothing of the data. Furthermore, we computed power spectra of Stokes $Q$ and $U$, $P$ and $RM$, both for the models and for the observations. The power spectrum of a field $X$ is defined as $PS_X(k) = |\mathcal{F}(X)|^2$, where $\mathcal{F}$ denotes a Fourier transformation, and $X$ can be $Q$, $U$, $P$ or $RM$. The power spectra of the static models are given as a function of wave number $k$, where $k = 1$ corresponds to the size of the cube (512 pixels). The power spectra of the dynamical model and the observations are given as a function of multipole $\ell$. Multipoles are similar to wave numbers, but are scaled to angular scales, viz. $\ell = 180^\circ/\theta$, where $\theta$ is the angular scale on the sky in degrees. The power spectra are computed in two dimensions, and azimuthally averaged.

9.3.1 The dynamical model

The $RM$ obtained from the dynamical model by integrating the line of sight component of the magnetic field times electron density over the path length, is given in Fig. 9.1. The $RM$ ranges from $-23$ to 19 rad m$^{-2}$ and has an average of $-0.3$ rad m$^{-2}$. Fig. 9.2 shows apparent $RM_x$ (top) and $P_z$ (bottom) derived from the smoothed $Q$ and $U$ maps, where the smoothing is 2, 4, and 8 pixels from left to right. The apparent
Figure 9.2: Maps of apparent rotation measure $R M_s$ (top), and of apparent polarized intensity $P_\perp$ (bottom). The maps are derived from the smoothed maps $Q_s$ and $U_s$, for smoothings of 2 (left), 4 (center) and 8 pixels (right).

$R M_s$ maps show structure on many scales, most conspicuously also on scales below the “beam size”, in the form of one-pixel wide ridges of anomalously high and low $R M$. The polarized intensity maps show mainly structure on the scale of the “beam size”. Narrow depolarization canals of one beam wide are visible, which are caused by beam depolarization (see Section 9.6.1).

Power spectra of $Q$, $U$, $P$ and $R M$ are shown in Fig. 9.3. Estimated errors in the power spectra given in the figure are purely from the azimuthal averaging of the 2-dimensional power spectra, no “observational” errors were added to the model. The steepening of the slope of the power spectra at small scales (high $\ell$) is due to the convolution of the $Q$ and $U$ maps. Therefore, we only make a linear fit to the power spectra at low $\ell$, for $\log(\ell) < 2.7$ and $\log(\ell) < 2.4$ for smoothings 2 and 4, respectively. The resolution at $\sigma = 8$ is so low that no reliable linear fit is possible. The values of the fitted slopes of the power spectra are given in Table 9.2. The slope of the power spectrum of $Q_s$ seems to be flatter than that of the power spectrum of $U_s$, but the resolution is too low to draw a conclusion.

9.3.2 The static models

The $R M$ maps of the static models are constructed from $Q$ and $U$ maps with added noise of 5%. Fig. 9.4 shows $R M$ maps for four static models $s1$, $s2$, $s3$, and $s4$ from left to right, and from top to bottom increasing smoothing; the top plot is the original $R M$, and below that smoothings of 2, 4, 6 and 8 are given. The grey scale saturates
Figure 9.3: Power spectra of Stokes $Q$ and $U$ (top), $P$ (center) and $RM$ (bottom) of the dynamical models. From left to right, smoothing is 2, 4, and 8. In the top plots, diamonds denote $Q$ and asterisks denote $U$. The solid lines are the linear fit made to the inertial range of the power spectrum. In the top plot, the linear fit to $Q$ is the solid line, the fit to $U$ the dotted line.

Table 9.2: Slopes $\alpha$ of the power spectra of $P$, $RM$, $Q$ and $U$ for dynamical and static models with smoothings $\sigma = 2, 4$, and 8, and for the observations. Errors in the models are only caused by the circular averaging of the 2-dimensional power spectra, and are $\sim 0.3$ for the dynamical model, and $\sim 0.1$ for the static models.

<table>
<thead>
<tr>
<th>$\sigma$</th>
<th>$\alpha_Q$</th>
<th>$\alpha_U$</th>
<th>$\alpha_P$</th>
<th>$\alpha_{RM}$</th>
</tr>
</thead>
<tbody>
<tr>
<td>$df$</td>
<td>0.9 0.8 -</td>
<td>1.6 2.0 -</td>
<td>1.6 2.2 -</td>
<td>1.8 1.6 -</td>
</tr>
<tr>
<td>$sl$</td>
<td>0.1 0.1 0.1</td>
<td>0.1 0.1 0.1</td>
<td>0.2 0.2 0.1</td>
<td>0.0 0.2 0.2</td>
</tr>
<tr>
<td>$s2$</td>
<td>0.1 0.2 0.3</td>
<td>0.0 0.1 0.0</td>
<td>0.1 0.1 0.1</td>
<td>0.2 0.2 0.4</td>
</tr>
<tr>
<td>$s3$</td>
<td>0.0 0.0 0.1</td>
<td>0.2 0.2 0.1</td>
<td>0.2 0.2 0.0</td>
<td>1.2 0.9 0.6</td>
</tr>
<tr>
<td>$sf$</td>
<td>0.3 0.3 0.2</td>
<td>0.2 0.3 0.3</td>
<td>0.4 0.5 0.4</td>
<td>1.8 1.4 1.3</td>
</tr>
<tr>
<td>$s5$</td>
<td>0.5 0.5 0.5</td>
<td>0.6 0.7 0.9</td>
<td>1.1 1.1 1.0</td>
<td>2.7 2.7 2.4</td>
</tr>
<tr>
<td>Aur</td>
<td>3.24 ± 0.14</td>
<td>3.42 ± 0.14</td>
<td>2.26 ± 0.25</td>
<td>0.99 ± 0.08</td>
</tr>
<tr>
<td>Hor</td>
<td>2.70 ± 0.14</td>
<td>2.85 ± 0.15</td>
<td>2.21 ± 0.32</td>
<td>0.94 ± 0.10</td>
</tr>
</tbody>
</table>
Figure 9.4: $RM$ maps for the static models $s1$, $s2$, $s3$, $s4$ (from left to right) in grey scale, which ranges from $-15 \text{ rad m}^{-2} < RM < 15 \text{ rad m}^{-2}$. From top to bottom: the original map derived from the models, $RM$ determined from smoothed $Q$ and $U$ maps with smoothing of 2, 4, 6, and 8. Regions of anomalous $RM$ grow with beam size. Note the structure introduced in the $RM$ distribution of model $s1$ which has a flat magnetic energy density power spectrum.

at $|RM| = 15 \text{ rad m}^{-2}$. Smoothing causes anomalously high or low $RM$s in patches of approximately the beam size, which is a different behavior from the $RM$ in the dynamical model. Maps of polarized intensity of the static models $s2$ and $s3$ are presented in Fig. 9.5. The smoothing applied to the $Q$ and $U$ maps is 4, and the grey
scale runs from 0 to 0.22. There is no large-scale structure visible in the $P$ maps.

Fig. 9.6 shows power spectra of $U$ and $Q$ in the same plot, and for $P$ for models $s1$ and $s5$. Smoothing increases from left to right as $\sigma = 2, 4$ and 8. Again, the decrease in power at large wave numbers $k$ is due to the smoothing, so the power spectrum can only be obtained for low wave number. The derived spectral indices are shown in Table 9.2. Errors result only from the azimuthal averaging of 2-dimensional power spectra. The power spectra of $Q$, $U$ and smoothed $P$ in the static models are almost flat, unlike the dynamical models and the observations (see Section 9.4.2).

To test the cause for the flat power spectra of $Q$, $U$ and $P$ in the static models, we irradiated the cubes of ISM of the static model $s2$ with a background polarized intensity with a power spectrum with multipole spectral index $\alpha = 2$, instead of the constant background polarized intensity used in the models. Fig. 9.7 shows the resulting power spectra of $P$, $Q$ and $U$ (no smoothing is applied). As expected, $P$ is unaffected by the Faraday screen, but $Q$ and $U$ flatten out above $k = 10$. This could be due to a number of differences between the static models and the observations, such as the lack of coupling between magnetic field $B$ and thermal electron density $n_e$, or no correlation of phases of spectral modes of the distributions of $B$ and $n_e$.

Power spectra of $RM$ of models $s2$ and $s5$ are shown in Fig. 9.8 for smoothings 2, 4, and 8. The $RM$ derived from the smoothed maps of $Q$ and $U$ are shown by the symbols, while the superimposed dotted line is the power spectrum of the original unsmoothed $RM$. The errors are in this case not only due to the azimuthal averaging of the power spectra, but also include a noise term in the $Q$ and $U$ maps of 5% of the $Q$ and $U$ signal. Again the spectral indices are given in Table 9.2.
9.4 The observations

9.4.1 Observational techniques

We compare the results of the models with observations of two regions, which are discussed in detail in Chapters 4 and 5. The first region, in the constellation of Auriga, is centered at \((l, b) = (161^\circ, 16^\circ)\). The second region, in the constellation Horologium, is centered at \((l, b) = (137^\circ, 7^\circ)\). The observational techniques used are discussed in Chapter 2, and are only briefly summarized here.

The fields were observed with the Westerbork Synthesis Radio Telescope (WSRT) in
Figure 9.7: Power spectra of $Q$ (diamonds), $U$ (triangles) and $P$ (squares) if the static cube $s2$ is irradiated with a background polarized intensity with power spectrum $PS(k) \propto k^{-2}$.

Figure 9.8: Power spectra of $RM$ in static models $s2$ (top) and $s5$ (bottom), for smoothings of 2, 4, and 8 from left to right. The symbols denote the $RM$ as derived from the smoothed $Q$ and $U$ maps, while the dotted line is the power spectrum of the original $RM$.

5 frequency bands of 5 MHz wide, centered at 341, 349, 355, 360 and 375 MHz. These frequencies correspond to the five wavelengths of 88 cm to 80 cm, for which $Q$ and $U$ maps were computed in the simulations. The observed regions are approximately 60 square degrees in size, and have a resolution of $\sim 5.0' \times 5.0'$ cosec $\delta$.

Because the minimum baseline length attainable with an interferometer is finite, an interferometer is insensitive to large scales. The WSRT is insensitive to scales larger than about $1^\circ$, which means that large-scale components in the $Q$ and $U$ maps are not detected. This could cause incorrect maps of polarization angle and therefore incorrect $RM$ determinations. However, we estimated that this effect is not important in these
two fields, see Chapter 3.

9.4.2 Results from the observations

Fig. 9.9 gives the polarized intensity at 349 MHz (left) and RM superimposed on $P$ (right) in the field in the constellation Auriga. $P$ and RM in the field in Horologium are shown in Fig. 9.10. In both fields, small-scale structure in polarized intensity $P$ and polarization angle $\phi$ is ubiquitous, but no small-scale structure in total intensity $I$ is detected. This indicates that the structure in polarization cannot be caused by structure in emission, but must be due to Faraday rotation and depolarization mechanisms.

Results from power spectra of $Q$, $U$, $P$ and RM in the observations are discussed in Chapter 8, and only the derived spectral indices are given in Table 9.2.

The rotation measures are obtained from the linear relation between $\phi$ and $\chi^2$. The polarization angle $\phi$ is determined up to $\pm 180^\circ$, but the observed range in RM is so small that this does not play a role in our observations (see Chapter 3). The relation between $\phi$ and $\chi^2$ is not always perfectly linear. This can be explained by depolarization mechanisms, which can destroy the linear $\phi(\chi^2)$-relation, and give an incorrect RM value. Therefore, we only consider “well-determined” RM values for which the reduced $\chi^2$ of the $\phi(\chi^2)$-relation $\chi^2_{\text{reduced}} < 2$, and the polarized intensity averaged over the frequency bands ($P > 20$ mJy beam ($\sim 4 \cdot 5$ times signal-to-noise).

From a depolarization model of the ISM which simulates the emission and propa-
gation of polarized radiation in a magneto-ionic medium, we find that a “polarization horizon” or “Faraday depth” can be derived (Chapter 3). Most of the polarized radiation originating at larger distances than the polarization horizon will be depolarized before it reaches the observer. Therefore, the average path length for polarized radiation at these low frequencies is estimated to be 600 pc, with a range of plausible values of 400 pc to 1200 pc.

9.5 Comparison of models and observations

In Fig. 9.11, we compare the distributions of $RM$ for the dynamical and static models with the observed ones. The histograms of the observations only include $RM$ values for which $\chi^2_{red} < 2$ and $\langle P \rangle > 20$ mJy/beam. The shoulder at $RM = -8$ rad m$^{-2}$ in the histogram of the Auriga field is due to the high intensity structure in the northwestern corner of the map, which has a relatively large negative $RM$. In Fig. 9.9, a large-scale gradient in $RM$ can be seen. If this $RM$ gradient is subtracted from the $RM$ map, the remaining $RM$ distribution is approximately Gaussian, like the distribution in the Horologium field.

The $RM$ distribution of the dynamical model is more or less Gaussian too, but the $RM$ histograms of the static models vary strongly with the steepness of the power spectrum, because a steep spectrum indicates much large-scale structure in $RM$, which destroys the Gaussian distribution. Therefore, judging from Fig. 9.11, the $RM$ distributions suggest that power spectra with a slope above $\alpha \approx 2.5$ are too steep to produce a Gaussian $RM$ distribution as is observed.

In the observations, the power spectra of $Q$ and $U$ are steeper than that of $P$, most likely due to a Faraday screen in front of the observations. Namely, a foreground Faraday screen induces additional structure on relatively large scales in $Q$ and $U$, but
not in $P$. In the simulations, power spectra of $Q$ and $U$ that are steeper than that of $P$ are not expected, because there is no large-scale foreground screen imprinting structure on a medium exhibiting smaller-scale structure. Instead, the spectral indices of $U$ and $P$ are similar to the spectral index of $P$ in the observations. The power spectrum of $Q$ appears to be flatter, but the resolution is too low to allow a robust conclusion.

In the static models, the input power spectrum of $B$ is known. Because the structure in $n_e$ has very low amplitude, the power spectrum of $RM$ is expected to resemble the power spectrum of $B$. This is not the case in models $s2$ and $s3$. The implication of this is not obvious, but it might be caused by the driver, which induces additional large-scale structure on the discrete scales of wave numbers $k = 1, 2$.

The static models are less realistic than the dynamical models, because of the lack of coupling between $B$ and $n_e$, and because they assume random phases of the spectral modes. As power spectra are squares of Fourier transforms, the Fourier phases do not influence the power spectra of $RM$ as long as the power spectra of $B$ and $n_e$ are isotropic. However, the spatial $RM$ distribution can be influenced by correlated phases of spectral modes, and therefore $Q$, $U$ and $P$ maps as well. Furthermore, in the presence of a uniform magnetic field, the isotropy assumption is violated.

The dynamical model is closer to reality in the sense that it includes coupling between $n_e$ and $B$, and that the evolution of the medium due to the driver allows for
phase correlations of Fourier modes. However, the resolution of the dynamical model is so low that numerical diffusion significantly influences the results. First, the steep spectrum of $RM$ in the dynamical simulation ($\alpha_{RM} \approx 2$ while in the observations $\alpha_{RM} \approx 1$) is probably caused by dissipation. Secondly, the largest scale on which the driving takes place is the scale of the short side of the bar-like domain, while the ISM is probably driven on larger scales (MacLow and Ossenkopf 2000), which e.g. can cause a large-scale component in the Galactic magnetic field. Furthermore, the variation in electron density in the dynamical model is higher than is allowed from the observations. This is probably due to the unrealistic velocity driver, which is not purely solenoidal, and therefore enhances electron density variations.

### 9.6 The influence of beam depolarization

In all radio polarimetric observations where structure in polarization is present on scales smaller than the telescope beam width, beam depolarization is important. However, in a medium that emits synchrotron radiation and Faraday-rotates, beam depolarization is not the only depolarization mechanism. Depth depolarization plays a rôle as well (see Section 9.1), so that it is difficult to estimate the influence of beam depolarization alone on the observations. The numerical simulations we discuss here do not include depth depolarization, so that beam depolarization is the only mechanism that decreases the polarized intensity. Therefore, the influence of beam depolarization on the $RM$ distribution can be studied for different scales of structure within the beam.

First of all, smoothing of the $Q$ and $U$ maps has a clear effect on the $RM$ maps, as can be seen from Figs. 9.2 and 9.4. The appearance of $RM$ is different for static and dynamical models. In the dynamical model, beam depolarization causes elongated one-dimensional “ridges” of anomalously high or low $RM$, with structure on scales smaller than the beam size. In these ridges, the $RM$ is badly determined, and the linear fit of $\phi$ versus $\lambda^2$ shows a high $\chi^2$. The ridges of anomalous $RM$ correspond to the depolarization canals in the $P$ maps. Due to the low $P$, polarization angle $\phi$ is badly determined, which causes anomalous $RM$ determinations. A $RM$ map of the observations in the Auriga field is shown in Fig. 9.12. In this figure, the $RM$ determinations from the linear fit of $\phi$ to $\lambda^2$ are taken at face value, i.e. including unreliably determined $RM$s with high $\chi^2$. The maps of apparent $RM$ in the dynamical model show the same behavior, i.e. narrow ridges of anomalous $RM$ values on scales smaller than the beam size at the positions of the depolarization canals. On the other hand, the static models show areas of anomalous $RM$ values as large as the “beam size”, which is not seen in the observations nor in the dynamical models.

The effect of beam depolarization on the determination of $RM$ can be quantified by the normalized difference between original $RM$ and the $RM$ from the smoothed maps $RM_s$ as

$$\delta RM = \sqrt{\langle (RM - RM_s)^2 \rangle / \langle RM^2 \rangle}.$$ 

where $\langle \rangle$ denotes averaging over the map. The $RM$ differences for the static and dynamical models are shown in Fig. 9.13 as a function of the amount of smoothing $\sigma$. Only well-determined $RM$s are included, which are defined as the $RM$s where the linear
Figure 9.12: Map of RM in the observations of the Auriga region, where RM values are shown for each pixel, regardless of quality of fit. The grey scale saturates at $|RM| = 15$ rad m$^{-2}$. The resolution is $\sim 5'$, so that the data is approximately 5 times oversampled. Structure in RM on scales smaller than the beam width is present in one-dimensional “ridges” of anomalous RM, at the positions of depolarization canals in the P map.

Figure 9.13: Normalized quadratic difference between original RM and RM$_e$ from smoothed Q and U maps for the static models (left) and the dynamical model (right). The diamonds denote model s1, triangles are s2, squares s3, crosses s4 and asterisks s5.

$\phi(\chi^2)$-fit has a reduced $\chi^2 < 2$ and $P > 5$ times signal-to-noise. Especially in the static models with steep power spectra ($\alpha \gtrsim 1$), even a smoothing $\sigma = 1$ instantaneously induces $\delta$RM$_e$s of up to 50% to 100%. In the dynamical model, all smoothing above $\sigma = 1$ results in a deviation of RM from the true values of more than 50%. So a RM
determination is uncertain at positions where the polarized intensity is significantly influenced by beam depolarization. However, the statistical distribution of $RM$ is less affected by the beam depolarization, as is shown in Fig. 9.14. This figure shows $RM$ distributions for the dynamical model, for the original $RM$ and apparent $RMs$ when smoothing the $Q$ and $U$ maps. The width of the $RM$ distribution remains similar, although peaks in the distribution are created by smoothings with $\sigma \geq 4$.

These results indicate that beam depolarization modifies individual $RM$ values, but that it does not significantly affect the statistical distribution of $RM$ for smoothing with $\sigma \leq 4$. Furthermore, $RM$ power spectra do not seem to be significantly influenced by beam depolarization (see Table 9.2), although the resolution of the dynamical model is too low to allow a strong conclusion.

Larger smoothing causes more depolarization. Therefore, comparison of the amount of depolarization in the simulations with the depolarization in the observations, allows us to estimate the amount of structure on sub-beam size scales that must be present in the observations. In Fig. 9.15 we present the amount of depolarization caused by smoothing for the 5 static models (left) and the dynamical model (right). The upper limit for the amount of beam depolarization in the observations is obtained by assuming that all depolarization in the observations is caused by beam depolarization. The average degree of polarization $\langle \rho \rangle = \langle P \rangle / \langle I \rangle$ in the Auriga region is about 7% and in the Horologium region 6%. With a maximum intrinsic degree of polarization of synchrotron radiation $\sim 70\%$ (Burn 1966), corresponding to $P = 1$ in the simulations, $\langle P \rangle$ in Fig. 9.15 is 0.1 and 0.09 for the Auriga and Horologium regions respectively. For the shallow spectra $s1$, $s2$ and $s3$, this indicates that $\sigma \leq 3$, and for the somewhat steeper spectrum $s4$ that $\sigma \leq 8$.

On the other hand, the dynamical model shows much less decrease in polarized intensity with increased smoothing. Fig. 9.15 indicates that structure on scales at least 10 times smaller than the beam is needed to explain the observed $P$ by beam depolarization alone. Structure in the ISM is believed to be present down to scales below an arcminute (Clegg et al. 1992). However, this cannot dominate in our fields because we observe coherent structure in polarization angle over many independent beams. This appears to confirm the conclusion in Chapter 3 that beam depolarization alone cannot be responsible for the observed structure in $P$. 

![Figure 9.14: Histograms of $RM$ for the dynamical model. From left to right the original $RM$, and $RM$ for smoothings $\sigma = 2$, 4, and 8.](image-url)
The steep decline of averaged polarized intensity with smoothing in the static models could be due to the lack of correlation between phases of the spectral modes. Absence of a phase correlation means minimum correlation between pixels. The static models show the maximum mean variation in $Q$ and $U$ between adjacent pixels, which results in more beam depolarization than for models with correlated Fourier phases. This indicates that models with random Fourier phases present an upper limit to the amount of beam depolarization.

### 9.6.1 Depolarization canals

Both static and dynamical models, as well as the observations, show narrow, onedimensional structures of low polarized intensity, which are “canals” of depolarization. Haverkorn et al. (2000, Chapter 6) showed that the average $\Delta \phi$ across canals in the Auriga region is $90^\circ (\pm 180^\circ)$. This is demonstrated in Fig. 6.3, which shows the distributions of the change in polarization angle $\Delta \phi$ over a telescope beam, measured as $\Delta \phi = \phi_i - \phi_j$ over two adjacent beams 1 and 2. In the top plots, $\Delta \phi$ was computed across a depolarization canal, whereas in the bottom plots, no canal was present in between beams 1 and 2. The distribution of $\Delta \phi$ across canine is centered at $90^\circ$, whereas the $\Delta \phi$ distribution excluding canals is centered around $0^\circ$, as expected. Haverkorn et al. argued that the canals are most likely created by beam depolarization. The canals indicate the boundary between two regions of fairly constant polarization angle $\phi_i$ and $\phi_j$, while the difference in angle between the two regions is $\Delta \phi = \phi_i - \phi_j = (n + 1/2) 180^\circ$, probably caused by an abrupt change in $RM$ (Chapter 3).

The same distributions of $\Delta \phi$ were computed for the static and dynamical models. The result is presented in Fig. 9.16 and is very similar to the observations. Only a smoothing of 2 is shown in this figure, but higher smoothings give the same results.

In the observations, other depolarization processes could play a role in the creation of the canals, although beam depolarization is probably the dominating process (see Chapter 3). In the simulations, beam depolarization is the only process which creates...
structure in polarized intensity. This unambiguously shows the effectiveness of beam depolarization in the creation of depolarization canals.

9.7 Conclusions

At the present stage of the project, detailed quantitative conclusions cannot yet be drawn about the comparison of the static and dynamical models with the observations. As the simulated medium does not contain emission and only acts as a Faraday screen, beam depolarization is the only depolarization mechanism in the models. The Galactic ISM is believed to be an emitting and Faraday-rotating medium, and therefore depolarization along the line of sight plays a rôle as well.

However, some preliminary conclusions can be given. Static models which exhibit a steep power spectrum of $B$ ($\alpha \geq 2.5$) fail to reproduce the approximately Gaussian distributions of $RM$ in the observations and the dynamical model. Polarized intensity maps in both types of models and in the observations show depolarization canals. In the observations, the canals are likely to be caused by beam depolarization, which is the only possible cause for the canals in the simulations. Maps of $RM$ produced from smoothed maps of $Q$ and $U$ in the dynamical model show a behavior similar to the observations, i.e. one-pixels wide “ridges” of anomalously high or low $RMs$ at the positions where depolarization canals are visible in the $P$ maps. From the above, we tentatively conclude that the dynamical model better approximates the observations than the static models. This is not unexpected, because the dynamical simulation allows for correlations between Fourier phases, and includes coupling between electron density and magnetic field. However, higher resolution is essential.

Beam depolarization destroys the linear relation between $\phi$ and $\lambda^2$ and therefore
MHD simulations of the warm ISM

yields incorrect values of $RM$. In both the dynamical and static models, the individual $RM$ values deviate by more than 50% even for moderate smoothing of $\sigma = 1$. However, the statistical distribution of $RM$ remains similar to the unsmoothed distribution for smoothings $\sigma \lesssim 4$.

Beam depolarization decreases the degree of polarization. In the static models with (relatively) shallow spectra ($\alpha \lesssim 2.5$), structure in polarization angle on scales of about 0.3 of the beam size is sufficient to explain the observed degree of polarization. However, the decrease of degree of polarization with smoothing is much less in the dynamical model, where structure on scales of less than 0.1 times the beam size is needed to reproduce the amount of depolarization in the observations. As in the observations structure in polarization is coherent over many beams, we believe that structure on such small scales does not dominate in the observations. This is an indication that depth depolarization plays an important rôle in the observations as well.

9.8 Future work

There are considerable opportunities for extending this work in the future, and we intend to pursue these. An attempt will be made to solve the problems that exist at the moment with the static and dynamical models, i.e. the cause of the flat $Q$, $U$ and $P$ power spectra in the static models, and the low resolution in the dynamical one. With improved models, the influence of beam depolarization on the $RM$ power spectrum and on the quality of the linear $\phi(\chi^2)$-fit can be estimated more reliably. This will enable quantitative comparison of the simulations and observations, with the eventual goal of being able to reliably derive the Galactic magnetic field power spectrum in the ISM from the observed $RM$.

In addition, a more sophisticated treatment of the radiation transfer process in the ISM needs to be made. The simulated medium acts as a Faraday screen, whereas the Galactic ISM is most likely both emitting and Faraday-rotating synchrotron radiation (see Chapter 3). Synchrotron emission in a magneto-ionic medium causes depth depolarization, which will alter the $Q$, $U$ and $P$ distributions. The $RM$ distribution will be changed in two ways by depth depolarization. First, depolarization will change polarization angles and yield badly determined $RM$. Secondly, an increasing amount of radiation is depolarized at increasing distances from the observer, effectively causing a Faraday depth, or polarization horizon. If we neglect the first effect, we can estimate the magnitude of the second by shortening the path length of the radiation in the simulations. It will be interesting to investigate depth depolarization in the simulations as well. Therefore, the propagation of synchrotron emission through the simulated medium needs to be considered properly.

Acknowledgements

We thank P. Katgert and E. G. Zweibel for fruitful discussions and suggestions to improve the paper. Computations presented here were performed on the SGI Origin 2000 machine of the Rechenzentrum Garching of the Max-Planck-Gesellschaft and on the SGI Origin 2000 cluster of the NCSA, University of Urbana-Champaign/Illinois.
MH is supported by NWO grant 614-21-006. FH is supported by a Feodor-Lynen grant of the Alexander von Humboldt Foundation and by (U.S.) National Science Foundation grants AST-9800616 and AST-00098701. The Westerbork Synthesis Radio Telescope is operated by the Netherlands Foundation for Research in Astronomy (ASTRON) with financial support from the Netherlands Organization for Scientific Research (NWO).

References

MacLow, M.-M., 2002, in *Simulations of magnetohydrodynamic turbulence in astrophysics*, ed. by T. Passot & E. Falgarone
Nederlandse samenvatting

Dit is een samenvatting van mijn proefschrift, geschreven voor niet-sterrenkundigen. Mijn promotieonderzoek is begeleid door dr. P. Katgert van Sterrewacht Leiden, en prof. dr. A. G. de Bruyn van Stichting ASTRON in Dwingeloo en het Kapteyn Instituut van Universiteit Groningen. De titel van het onderzoek is "WSRT Polarimetrie van het warme ISM". Hieronder hoop ik duidelijk te maken wat dit inhoudt.

1 U bevindt zich hier

Eerst een korte schets van de plaats van onze zon in het heelal, en onze omgeving (op astronomische schalen), ter referentie. “De Melkweg” is vooral bekend als band van vaag wit licht aan de hemel, te zien bij heldere nachten. Dit licht komt van duizenden sterren die te zwak zijn om individueel waar te nemen. Deze sterren vormen ons sterrenstelsel, de Melkweg, bestaande uit vele miljarden sterren waaronder onze zon.

De Melkweg is een platte, ronddraaiende schijf met spiraalarmen, met een diameter van ongeveer honderdduizend lichtjaar. Het heelal is gevuld met miljarden sterrenstelsels: spiraalstelsels zoals het onze, elliptische en onregelmatige stelsels, elk met miljoenen tot miljarden sterren erin.

Fig. 1 geeft aan hoe de Melkweg er van grote afstand uit ziet. In het Melkwegcentrum bevindt zich de kern. Hierin, en in de spiraalarmen, zijn de meeste sterren geconcentreerd. Zoals in de figuur te zien is, staat de zon zo'n 30.000 lichtjaar verwijderd van het centrum, net aan de buitenkant van een spiraalarm.

2 De ruimte tussen de sterren

Als we op een heldere donkere nacht naar boven kijken, zijn sterren het enige wat we met het blote oog van de Melkweg kunnen zien. Telescopen bieden niet alleen de mogelijkheid om scherper en meer te kunnen zien dan met het blote oog, maar kunnen ook andere soorten straling zichtbaar maken, zie sectie 4.1. Dan blijkt wat zich nog allemaal tussendoors bevindt: gigantische gas- en stofwolken, resten van gestorven sterren, materie die uitgebrand is door sterren, supersnel bewegende deeltjes genaamd kosmische straling enzovoort. Alles wat zich tussen de sterren bevindt wordt het interstellair medium (ISM) genoemd, en is het onderwerp van onderzoek van veel astronomen. Een groot deel van het ISM bestaat uit gas in drie toestanden: koud gas (met een temperatuur van zo'n -200° C), warm gas (10.000° C) en heet gas (1.000.000° C). Het warme gas tussen de sterren is het onderwerp van dit proefschrift. Maar waarom zouden we ons druk maken over gas wat we niet eens kunnen zien?

1Eén lichtjaar is de afstand die het licht in een jaar aflegt, nl. 9.460.730.500.000 km
2Dit is geen echte foto van de Melkweg. Immers, we zitten zelf middenin het stelsel, dat veel te groot is om eruit te reizen en een overzichtsfoto te maken. In werkelijkheid is dit een foto van het nabije spiraalstelsel NGC 1232, dat erg op de Melkweg lijkt.
3Met "gas" bedoel ik gasvormige materie, dus bestaand uit losse moleculen en atomen, zoals bijvoorbeeld lucht (al bestaat het gas tussen de sterren uit andere stoffen dan lucht).
Figuur 1: Een "bovenaanzicht" van de Melkweg. De Melkweg is een platte schijf met spiraalarmen en een kern, waarin de meeste sterren zich bevinden. De zon ligt aan de rand van de Melkweg, ongeveer 30.000 lichtjaar van de kern, net aan de buitenkant van een spiraalarm.

2.1 Kraamkamer en kerkhof

Een van de redenen waarom het gas tussen de sterren een belangrijke rol speelt in de Melkweg is dat er sterren van gemaakt worden. Als een gaswolk maar genoeg samenbrekt, wordt die wolk uiteindelijk zo dicht dat er een ster kan ontstaan. Die ster zondt straling uit, die wisselwerkt met het interstellaire gas. Sommige sterren hebben ook krachtige sterrenwinden van gas en stof, die vanaf het oppervlak van de ster weggeblazen worden. Sterren zelf bewegen ook ten opzichte van het gas. (Dat desondanks de posities van sterren niet lijken te veranderen komt alleen doordat ze zo ver weg staan.) Ook de zon, met alle planeten die eromheen draaien, beweegt door een gaswolk heen, die we met telescoop kunnen waarnemen.

Als een ster uiteindelijk sterft, stoot die zijn buitenlagen weer uit, terug de interstellaire gaswolken in. Het interstellaire medium is dus zowel de kraamkamer als het kerkhof van sterren, en is een essentiële schakel in de kringloop van stellaire geboorte en dood. Onze zon, en dus ook de aarde, is ook ontstaan uit deze gaswolken. Zo bestaan ook wij uit interstellaire materie, en hebben al onze atomen al een of meerdere keren een ster van binnen gezien!
2.2 Wat gaswolken beweegt

De gas- en stofwolken zijn geen statische, inerte wolken zoals lome stapelwolken op een warme zomerdag, maar zijn dynamische, bewegende structuren, zoals donderwolken tijdens een goede storm. Verschillende processen veroorzaken beweging van het gas. Ten eerste kunnen zware sterren (meer dan drie keer zo zwaar als onze zon) gigantische explosies geven als ze sterven, zoogenaamde supernovae. Door een supernova worden de buitenlagen van de ster samen met het omringende gas alle kanten uit geblazen. Schokgolven kunnen dan door het gas gaan lopen, zoals een boeggolf door het water loopt als er een schip langskomt en op grote afstanden van het schip nog schade aan kan richten.

Een andere factor die heel belangrijk is in de bewegingen van het interstellair gas zijn de magnetische velden die in het gas aanwezig zijn.

3 Magneetvelden in de ruimte

Bekende magneten op aarde variëren van koelkastmagneetjes tot het aardmagnetisch veld, maar het enige wat we in het dagelijks leven van magneetvelden merken is dat ze voorwerpen laten plakken aan de koelkast en dat kompassen naar het noorden wijzen. In de interstellaire ruimte is dit totaal anders, daar is de invloed van magneetvelden overal merkbaar. Eerst zal ik uitleggen wat magneetvelden zijn, en hoe ze werken in de ruimte.

Een magneetveld is de manier waarop we de aantrekkingskracht van magneten beschrijven. Die aantrekkingskracht is tot grote afstanden van een magneet aanwezig, hoewel de kracht steeds zwakker wordt als de afstand tot de magneet groter wordt. Een magneetveld wordt weergegeven met magnetische veldlijnen, zoals getekend in Fig. 2. De lijnen met pijlen geven aan in welke richting het magneetveld staat, m.a.w. waar een kompasnaald naar toe zou wijzen als het kompas op een veldlijn geplaatst zou worden. Bovendien staan de lijnen dichter bij elkaar waar het magneetveld sterker is. De linkerafbeelding in Fig. 2 geeft het simpele magneetveld van een staafmagneet weer, dat het sterkst is bij de uiteinden (noord- en zuidpool) van de magneet. Daar staan dan de magneetveldlijnen het dichtst bij elkaar getekend. Verder weg van de magneet staan de veldlijnen verder uit elkaar, wat betekent dat de aantrekkingskracht daar zwakker is. Iedereen die wel eens gespeeld heeft met een staafmagneet en een papierclip weet dit ook uit eigen ervaring. Het magneetveld van de staafmagneet is een simpel en regelmatig magneetveld, met geordende magneetveldlijnen. Zoals de rechterfiguur laat zien, hoeft dit helemaal niet het geval te zijn. Dit is een plaatje van de zon (in het wit, met een asymmetrische zonnewind eromheen). In de figuur zijn magneetische veldlijnen getekend om te laten zien hoe het zonnemagneetveld loopt. Dit is totaal chaotisch, met structuren op veel verschillende schalen, en kan in korte tijd (uren tot weken) compleet veranderen.

De zon, en ook het interstellaire gas, heeft zo'n chaotisch magneetveld, doordat dit veld een sterke wisselwerking heeft met het gas. Het gas in de zon en in een groot deel van het interstellaire medium is namelijk geen gewoon gas zoals we dat op aarde kennen, maar een geloei-gerookte gas, ofwel een plasma. Een van de eigenschappen van plasma is dat het “vastkleeft” aan een magneetveld, zodat de magnetische veldlijnen
als het ware vastgeweven zijn in het gas. Dus als een gaswolk wil bewegen, kan deze tegengehouden worden door het magneetveld, als dat sterk genoeg is. Is de beweging van de wolk sterker dan het magneetveld, dan zal het veld worden meegesleurd door de gaswolk. Hierdoor heeft het magneetveld, hoewel onzichtbaar, toch een grote invloed op het interstellaire gas, en daarmee op de vorming van sterren.

Deze grote invloed van het magneetveld in de Melkweg op het gas en de sterren maakt het een interessant stofobject. Dit werpt meeteen een probleem op: hoe kun je het dan bestuderen? Magneetvelden zelf zijn niet te zien, met welke telescoop dan ook, dus die kunnen alleen indirect waargenomen worden. In dit proefschrift beschrijf ik metingen van straling die op een bepaalde manier veranderd is door het magneetveld in de Melkweg. Uit deze verandering van de straling bereken ik wat voor soort magneetveld daarvoor verantwoordelijk is.

4 Gepolariseerde radiostraling

De straling die ik in dit proefschrift heb onderzocht is gepolariseerde radiostraling. Ik zal daarom eerst uitleggen wat radiostraling is, wat het begrip polarisatie inhoudt, en hoe gepolariseerde radiostraling gemeten wordt.

4.1 Wat is radiostraling?

Radiostraling is net als zichtbaar licht, Röntgenstraling, infraroodstraling, microgolfstraling etc. *elektromagnetische straling*. Deze straling bestaat uit golven die zich door de lucht (of de ruimte) voortbewegen. Het enige verschil tussen de soorten straling is de golflengte, dus de lengte die één golf van die straling heeft, zie Fig. 3. Zichtbaar licht is elektromagnetische straling met een golflengte van ongeveer een miljoenste millimeter, terwijl bijvoorbeeld Röntgenstraling een nog veel kleinere golflengte heeft. Infraroodstraling (wat wij waarnemen als warmte) heeft een iets langere golflengte dan zichtbaar
licht. Microgolven (die een magnetron uitzendt) zijn nog wat langer, en radiostraling heeft de langste golflengte, nl. van centimeters tot kilometers lang.

Iedereen weet dat licht wordt tegengehouden door een plaat karton, en warmte door een dikke muur. Maar radiostraling laat zich door bijna niets tegenhouden. Daarom werken radio-, tv- en gsm-antennes ook met radiostraling: als er toevallig een flatgebouw tussen een zendmast en een antenne staat, gaat de straling daar gewoon doorheen, en kan er (bijna) overal televisie gekeken worden. Dit is een van de redenen waarom radiostraling ook uitermate geschikt is om astronominische waarnemingen te doen. Door dat het niet tegengehouden wordt door gas- en stofwolken, kun je met radiotelescopen heel ver weg kijken, veel verder dan bijvoorbeeld met optische telescopen voor zichtbaar licht.

4.2 Hoe ontstaat radiostraling?

Veel verschillende objecten in ons heelal zenden radiostraling uit. Naast radio’s, televisies en gsm’s zijn dat bijvoorbeeld ver weg gelegen sterrenstelsels, ontploffende sterren en kosmische stralingsdeeltjes. Deze laatste zijn deeltjes (vooral protonen en elektronen) met een snelheid bijna zo hoog als de lichtsnelheid, die overal in de Melkweg in elke richting bewegen. Over hun oorsprong en de mechanismen verantwoordelijk voor hun verspreiding bestaat nog discussie. In interactie met het magneetveld in de Melkweg zenden ze radiostraling uit, die zich vervolgens door de Melkweg heen beweegt. Dit vormt als het ware een achtergrond van radiostraling die overal aanwezig is. Deze straling onderzoek ik in dit proefschrift. Het magneetveld in de Melkweg zorgt niet alleen voor het ontstaan van deze straling, maar modificereert de straling ook tijdens zijn reis door de Melkweg naar de radiotelescoop. Deze modificatie kunnen we waarnemen, en uit de manier waarop de radiostraling veranderd wordt, leiden we eigenschappen van het magnetische veld af.

4.3 Waarnemen met radiotelescopen

Radiostraling wordt waargenomen met radiotelescopen, die groot moeten zijn doordat de golflengte van radiostraling zo groot is. De telescoop die ik gebruik heb voor mijn onderzoek is de Westerbork Synthesis Radio Telescope (WSRT) in Westerbork, Drenthe. De telescoop bestaat uit 14 afzonderlijke telescopen, elk met een schotel van...
voortplantingsrichting van de golf

diagram van de polarisatierichting van de golf

**Figuur 4:** Links staan twee golven getekend, die zich beide van links naar rechts voortbewegen. De polarisatierichtingen van de golven zijn verschillend: verticaal voor de bovenste golf, horizontaal voor de onderste golf. De standaardmanier om polarisatierichtingen weer te geven staat rechts. De cirkel geeft een stralingsbundel weer die recht op de lezer afkomt, terwijl de pijlen erin weergeven in welke richting de straling gepolariseerd is.

25m in diameter, in een rij opgesteld met een tussenafstand die verschilt van 36 meter tot 1.4 km en met een totale lengte van 2.7 km.

Hoe groter een telescoop, hoe scherper je ermee kunt zijn, zoals ook een dikke bril meer vergrt dan een dunne bril. Nu is er in de radiosterrenkunde een techniek ontwikkeld, genaamd *interferometrie*, waarbij de straling die binnenvalt in de 14 telescopen gebundeld kan worden alsof de straling met één grote telescoop van 2.7 km groot gemeten wordt. Hiermee kan dus een grote telescoop gesimuleerd worden door een stel kleine op een rij te zetten.

### 4.4 Wat is polarisatie?

De polarisatie van een stralingsgolf is de *golfrichting* van die golf. Niet de richting waarin de golf zich voortbeweegt, maar de richting waarin de golf "golft". Dit is aangegeven in Fig. 4. Hier staan twee stralingsgolven geschetst, die zich allebei van links naar rechts bewegen. Echter, de golfrichting, ofwel polarisatierichting, van de bovenste golf is van boven naar beneden, en die van de onderste golf is van voren naar achteren. Aan de rechterkant van de figuur staat de manier waarop de polarisatie weergegeven wordt.

Een *lineair gepolariseerde* bundel van radiostraling is een bundel radiostraling waarvan alle golven in dezelfde richting golven. Als alle golven in de bundel in willekeurige richtingen golven, dan is de bundel *ongepolariseerd*. Het meeste licht wat we dagelijks zien, zoals bijvoorbeeld zonlicht en lamplicht, is ongepolariseerd. Als bijvoorbeeld de helft van alle golven in een bundel in een bepaalde richting golven, en de andere helft heeft een willekeurige golfrichting, dan spreken we van *partiëel gepolariseerde* straling, in dit geval 50% gepolariseerd. De straling heeft dan een zogenaamde *polarisatiedraad* van 50%.
5 Faraday draaiing van geëpolariseerde radiostraling

Faraday draaiing is een proces dat optreedt in het warme interstellaire gas, in aanwezigheid van een magneetveld. Het veroorzaakt een draaiing van de polarisatierichting van geëpolariseerde straling terwijl de straling door het interstellaire gas reist. Dus hoe langer de afgelegde afstand van de straling door het warme gas is, hoe meer de polarisatierichting van de straling gedraaid wordt.

De hoeveelheid Faraday draaiing in het gas is te meten door waarnemingen te doen op verschillende golflengtes die dicht bij elkaar liggen. De Faraday draaiing van de polarisatierichting verandert namelijk met golflengte. En hoe de draaiing verandert, hangt weer af van het magneetveld. Een voorbeeld hiervan wordt gegeven in Fig. 6. Deze figuur beschrijft waarnemingen over verschillende golflengtes λi t/m λf, en laat zien hoe de polarisatierichting verandert in een sterk en in een zwak magneetveld.

De maat voor de hoeveelheid Faraday draaiing is de rotatiemaat RM. De RM hangt af van drie grootheden: (1) de sterkte van het magneetveld, zoals boven al beschreven; (2) de dichtheid van het interstellaire gas: bij een hogere dichtheid is de draaiing sterker; en (3) de afstand die de straling door het warme gas aflegt (de padlengte): over langere afstanden kan meer draaiing opgebouwd worden.

Met de Westerbork telescoop is het mogelijk om de binnenkomende geëpolariseerde straling tegelijkertijd op verschillende golflengtes te meten. Uit de waargenomen draaiing van de polarisatierichting van die straling in metingen op verschillende golflengtes kan dan de RM afgeleid worden. Om uit de gevonden RM waarden informatie over het interstellaire magneetveld te vinden, moeten we een bepaalde dichtheid van het gas...
sterk magneetveld

\[
\begin{array}{cccc}
\uparrow & \rightarrow & \uparrow & \rightarrow \\
\end{array}
\]

golf lengtes: \( \lambda_1 \), \( \lambda_2 \), \( \lambda_3 \), \( \lambda_4 \)

zwak magneetveld

\[
\begin{array}{cccc}
\uparrow & \rightarrow & \uparrow & \rightarrow \\
\end{array}
\]

Figuur 6: Voorbeeld van Faraday draaiing. De vier cirkels naast elkaar stellen dezelfde bundel gepolariseerde straling voor, maar waargenomen op vier verschillende golf lengtes genaamd \( \lambda_1 \) t/m \( \lambda_4 \). Faraday draaiing van de polarisatierichting is verschillend per golflengte. De bovenste situatie geeft een sterk magneetveld in het warme gas weer, waarin de straling veel gedraaid wordt. In de onderste situatie is het magneetveld zwakker, wat een kleinere rotatie tot gevolg heeft.

en een bepaalde padlengte van de straling aannemen. Hoewel we hier wel schattingen van kunnen maken, zijn de precieze waarden onzeker.

6 Polarisatie en depolarisatie

Het tweede proces dat de radiostraling verandert is depolarisatie. Dit is het proces dat gepolariseerde straling verandert in ongepolariseerde straling, doordat twee gepolariseerde golven die een verschillende polarisatierichting hebben, elkaar (gedeeltelijk) uitdoven, zie Fig. 7.

Het volume interstellair gas waar gepolariseerde radiostraling wordt uitgezonden, is duizenden lichtjaren lang. De opgevangen straling is dan ook een opstelling van straling die op elke positie langs de gezichtlijn wordt uitgezonden. Dit optellen van gepolariseerde radiostraling kan depolarisatie veroorzaken. Als alle gepolariseerde golven (straling) die we in een bepaalde richting waarnemen dezelfde polarisatierichting hebben, zal er geen depolarisatie optreden, terwijl er bij wisselende polarisatierichtingen langs de gezichtlijn veel depolarisatie optreedt.

Een van de belangrijkste factoren die de polarisatierichting van straling beïnvloeden is het interstellaire magneetveld. Het magneetveld is heel complex, maar we kunnen het beschrijven door middel van twee componenten: een grote-schaal component en een kleine-schaal component. Het grote-schaal magneetveld is de component die heel geordend is over de hele Melkweg, met magnetische veldlijnen op regelmatige afstanden van
elkaar, en gericht langs de spiraalarmen. Met de kleine-schaal component wordt dan het chaotische magneetveld bedoeld, dat over korte afstanden verandert. In werkelijkheid is het magneetveld niet te scheiden in grote- en kleine-schaal componenten, doordat er op iedere schaal structuur aanwezig kan zijn. Echter, voor deze analyse is deze beschrijving een goede benadering. Welke component van het magneetveld het beste te zien is, hangt af van de methode van waarnemen, zoals een bakstenen muur glad en regelmatig lijkt als je die van ver bekijkt, maar bobbelig blijkt (kleinschalige structuur bevat) bij inspectie van dichtbij. We weten zeker dat beide componenten aanwezig zijn in het interstellair gas, maar wat de verhouding is van de grootschalige ten opzichte van de kleinschalige component, is (nog) niet precies bekend.

Als het magneetveld in het interstellair gas chaotisch is en veel kleine-schaal structuur bevat, is de polarisatiegraad laag en zien we weinig polarisatie. Terwijl een regelmatig magneetveld bijna niet depolariseert, zodat we in dit geval veel polarisatie zien. Dit is een heel belangrijk punt, wat betekent dat we uit de verschillen in hoeveelheid polarisatie kunnen afleiden hoe het magneetveld zich gedraagt!

Dit wordt geïllustreerd in Fig. 8. Hier worden drie voorbeelden gegeven van verschillende magneetvelden (rechts). De linkerfiguur laat zien hoe de polarisatie zich gedraagt in deze magneetvelden, en onderaan staat weergegeven hoe deze polarisatie dan waargenomen zou worden.

7 De waarnemingen

Het onderzoek beschreven in dit proefschrift heb ik natuurlijk niet alleen uitgevoerd, maar met mijn begeleiders en verschillende andere medewerkers. Daarom zal ik bij de beschrijving van het onderzoek de meervoudsvorm “we” gebruiken, en medewerkers buiten mijn eerdere genoemde begeleiders vermelden.

We hebben de Westerbork telescoop gebruikt om afbeeldingen te maken (“imaging”) van de gepolariseerde radiostraling, zogenaamde *polarimetrie*. Twee gebieden aan de hemel hebben we waargenomen in vijf verschillende golflengtes tussen ongeveer 80 en 88 cm. Deze twee gebieden zijn vrij groot, ongeveer 50 vierkante graden, wat overeenkomt met een vlak waarin ongeveer 200 volle manen passen. Het ene gebied

![Diagram](image-url)
bevindt zich in het sterrenbeeld Voerman (Latijnse naam: Auriga), en is het “Auriga veld” gedoopt; het andere gebied is in de richting van het sterrenbeeld Slingeruurwerk (Horologium) en heet dus het “Horologium veld”. De afbeeldingen van gepolariseerde radiostraling in de twee velden staan in Fig. 9. Deze kaarten laten zien welke richting veel gepolariseerde radiostraling (wit), en waar weinig gepolariseerde radiostraling (zwart) vandaan komt. In het Auriga veld lopen een paar rechte filamenten waar veel gepolariseerde radiostraling zichtbaar is, terwijl we in het Horologium veld veel gepolariseerde radiostraling opvangen uit een circulaire structuur. (Van die circulaire structuur is in Fig. 9 eigenlijk alleen de linkerkaart zichtbaar, als een boog links van het midden.)

In de jaren negentig is een groot project uitgevoerd met de Westerbork telescoop, waarin de hele noordelijke hemel werd waargenomen in radiostraling. Hierbij werd ook het gepolariseerde deel van de straling apart waargenomen en de resultaten opgeslagen, maar tot nu toe was er geen tijd om daarnaar te kijken. Nu verwerken en analyseren we deze data in samenwerking met Dominic Schnitzeler, een student in Leiden die op dit project zal afstuderen. Omdat zo'n groot deel van de hemel is waargenomen, is deze data ideaal om de polarisatie-eigenschappen in verschillende posities aan de hemel met elkaar te vergelijken.
Figuur 9: Kaarten van het gepolariseerde deel van de radiostraling in de twee waargenomen gebieden aan de hemel. Links het “Auriga veld”, rechts het “Horologium veld”. De golfomvang van de straling is 86 cm. Dit betekent dat er veel gepolariseerde straling uit die richting komt.

7.1 Faraday draaiing in de waarnemingen

We hebben rotatiematen $R_M$ afgeleid voor de twee waargenomen gebieden, zoals weergegeven in Fig. 10. Het eerste wat opvalt aan de verdeling van $R_M$s is hoeveel structuur er aanwezig is op kleine schalen. Die structuur komt waarschijnlijk door variaties in het magneetveld en in de dichtheid van het gas, doordat het gas gekoppeld is aan het magneetveld (zie sectie 3). Maar een verandering van het teken van de rotatiemoment (van negatief naar positief of vice versa) impliceert altijd een verandering van richting van het magneetveld. Analyse van deze rotatiematen geeft dus informatie over het interstellaire magneetveld, rekening houdend met de dichtheid van het gas, en de padlengte van de straling door het gas.

7.2 Depolarisatie in de waarnemingen

Doordat op sommige plaatsen veel gepolariseerde radiostraling waargenomen wordt, en op andere weinig, zou de eerste gedachte zijn dat er soms veel en soms weinig straling uitgezonden wordt. Maar, met de gepolariseerde straling zien we maar een deel van de totale (gepolariseerde en ongepolariseerde) straling. Bestudering van de totale straling die uitgezonden wordt, laat zien dat de totaal uitgezonden straling op alle plaatsen gelijk is! En deze straling wordt in het gas uitgezonden met een vaste polarisatiegraad. De fysische theorie van voortplanting van kosmische straling in een magneetveld dicteert namelijk dat de straling op het moment van uitzending ongeveer 70% gepolariseerd is. Dit betekent dus dat in de richtingen waarin we weinig gepolariseerde straling kunnen zien, de straling is gedepolariseerd.
Figuur 10: Rotatiemaat $RM$ voor het Auriga veld (links) en het Horologium veld (rechts). De grijschaal is hetzelfde als Fig. 9, en de cirkels zijn de rotatiematen. Doordat niet op elke positie een “goede” $RM$ afgeleid kon worden, staan niet overal $RM$s weergegeven. De grootte van de cirkel is geschaald met de grootte van $RM$. Open cirkels geven negatieve $RM$s aan, gevulde cirkels positieve.

Wat betekent dit nu concreet voor de waarnemingen die we hebben? Doordat een grote hoeveelheid gepolariseerde straling een regelmatig magnetenveld aanduidt, moeten in het Auriga veld (links) filamenten van heel regelmatig magnetenveld voorkomen. Of misschien zijn het vlakken met een regelmatig magnetenveld, waar we precies aan de zijkant tegenaan kijken. In het Horologium veld (rechts) moet er een magnetische structuur met een cirkelvorm zijn, en het meest waarschijnlijk is dat het een soort tunnel van magnetenveld is, waar we recht in kijken.

Nu klinkt dat op het eerste gezicht erg toevallig, een vlak waar we van de zijkant tegenaan kijken, een tunnel die we precies in de lengterichting zien: waarom zouden er dan ook niet vlakken en tunnels met andere oriëntaties zijn? Nu zijn die er zeer waarschijnlijk ook, maar we kunnen ze niet zien. Als je een vlak van de zijkant bekijkt, wordt een groot deel van de gezichtslijn ingenomen door dat vlak, terwijl als je het vlak recht van voren bekijkt, je maar door een klein stuk van het vlak kijkt. Hetzelfde geldt voor een magnetische tunnel. Een vlak of tunnel in een andere oriëntatie maakt dan een klein gedeelte van de gezichtslijn uit. De polarisatie wordt maar langzaam opgebouwd langs de gezichtslijn, waardoor een “kort” stuk gas met regelmatig magnetenveld niet voldoende is om genoeg polarisatie op te bouwen om de structuur waar te kunnen nemen. Dus alleen met bijzondere situaties, zoals een vlak van de zijkant of een tunnel in de lengterichting, kunnen we waargenomen structuren verklaard worden. Alle andere situaties leveren gewoonweg te weinig polarisatie op.
8 Computermodellen

8.1 Modellering van depolarisatie

Tot nu toe hebben we depolarisatie alleen gebruikt om het magneetveld in hele specifieke structuren zoals filamenten en de ring-structuur te schatten. Maar voor de studie van het typische magneetveld in het warme interstellaire gas, moet de typische depolarisatie over de waargenomen velden bekeken worden. Met dit doel hebben we een simpel computermodel ontworpen dat de depolarisatie in het interstellaire gas kan nabootsen. De situatie is zo complex dat vereenvoudigingen gemaakt moeten worden in het model. De belangrijkste vereenvoudiging is de aannemer dat de gaslaag kan worden beschreven als opgebouwd uit cellen met een bepaalde grootte, waarin zich gas bevindt met een bepaald magneetveld. We nemen weer aan dat het magneetveld bestaat uit een kleine-schaal component die verandert per cel, en een grote-schaal component die overal hetzelfde is. We berekenen de geop polariseerde radiostraling die uitgezonden wordt in het gemodelleerde gas, en hoeveel van die straling gedepolariseerd wordt terwijl de straling zich door gas voortbeweegt. De uitkomsten daarvan worden met de waarnemingen vergeleken. Door het magneetveld (grote- en kleine-schaal componenten) en de celgrootte in het model te variëren, kunnen we het magneetveld vinden waarmee de gemodelleerde polarisatie zoveel mogelijk op de waargenomen polarisatie lijkt, en dit voor verschillende celgroottes. In dat geval is het waarschijnlijk dat het magneetveld dat als input in het model gebruikt is, ongeveer het echte magneetveld in het interstellaire gas weergeeft. Het blijkt dat maar een beperkt aantal celgroottes toegestaan is bij vergelijking van het model met de waarnemingen, wat een bepaald bereik in toegestaan magneetveld oplevert.

Een van de belangrijkste uitkomsten van het model is dat het grote-schaal magneetveld sterker moet zijn dan het kleine-schaal veld. En dat is niet wat andere waarnemingen vinden: over het algemeen is in het interstellaire gas het kleine-schaal veld sterker dan het grote-schaal veld en niet omgekeerd. Echter, deze afwijkende waarde wordt heel logisch als rekening gehouden wordt met de richting waarin de velden waargenomen zijn. De richtingen waarin de Auriga en Horologium velden staan is de Melkweg uit (naar rechts boven in Fig. 1). Waarschijnlijk is de straling die in de telescoop wordt opgevangen, uitgezonden vanaf elke positie langs de gezichtlijn. Echter, als de straling op grote afstand wordt uitgezonden, is de kans groot dat deze voor het bereiken van de telescoop gedepolariseerd wordt. In de praktijk betekent dit dat de straling uit de ver weg staande spiraalarm in de waarnemingsrichtingen compleet gedepolariseerd is voordat het ons bereikt. Dus die spiraalarm is niet zichtbaar in gepolariseerde straling, en de straling die waargenomen wordt is alleen afkomstig van het gas tussen de spiraalarmen. De spiraalarmen bevatten de meeste sterren, en zijn dan ook de plaatsen van de meeste actie: de meeste sterrondballen, schokgolven en sterrenwinden vinden hier plaats. Doordat het magneetveld hierin meegesleurd wordt, verwachten we ook dat het magneetveld het meest chaotische is in de spiraalarmen, en een stuk regelmatiger daarbuiten. En dit is precies wat onze waarnemingen en modellen ons laten zien!
8.2 Modellering van het interstellaire medium

Het computermodel zoals boven beschreven is erg eenvoudig, en veel simplificeringen zijn gemaakt om de depolarisatie te beschrijven. Een medium zoals boven beschreven is zelfs zo simpel dat het in werkelijkheid niet eens kan bestaan. Daarom is dit alleen een speelgoedmodel ("toy model"), waarmee we ongeveer kunnen schatten hoe het interstellaire magneetveld zich gedraagt, maar dat geen gedetailleerde informatie kan geven.

Nu bestaan er zeer geavanceerde computermodellen van het interstellaire medium, waarin realistische gasstromingen en magneetvelden verwerkt zijn. Om dit soort modellen te kunnen gebruiken om onze waarnemingen te verklaren werk ik samen met Fabian Heitsch, een specialist in computermodellen van het interstellaire gas aan de Universiteit van Colorado te Boulder (VS). We berekenen hoe gepolariseerde radiostraling zich zou gedragen als het door het interstellaire gas van zijn modellen zou reizen, en vergelijken dat met de waargenomen gepolariseerde radiostraling. De verschillen en overeenkomsten tussen computermodel en waarneming geven ons vanuit een andere invalshoek informatie over het interstellaire magneetveld.

9 Conclusie

Magneetvelden in de interstellaire ruimte spelen een essentiële rol bij bijvoorbeeld ster-vorming, of de evolutie van de Melkweg. Kennis over interstellaire magneetvelden is daardoor noodzakelijk voor een goed begrip van het ontstaan en evolutie van onze Melkweg en specifieker van sterren, waaronder onze zon. Magneetvelden zijn echter moeilijk waar te nemen. Eén enkele methode om alle facetten van het magneetveld te kunnen vinden bestaat zelfs niet: er zijn verschillende methoden waarmee een component van het magneetveld, of het veld in een specifieke soort gas (bijvoorbeeld het warme gas) bestudeerd kan worden.

De methode van het waarnemen van de (de-)polarisatie en Faraday draaiing van radiostraling in het interstellaire gas is gebruikt in dit proefschrift. Dit is een vrij nieuwe methode, waarvan tot voor kort niet bekend was of deze echt gebruikt kon worden om nieuwe informatie over het interstellaire magneetveld in de Melkweg te vinden. Dit proefschrift toont aan dat dit werkelijk mogelijk is. Bijvoorbeeld, door de waarnemingen te combineren met modellen (zie sectie 8), kunnen we een schatting geven over hoe sterk kleine-schaal en grote-schaal componenten van het magneetveld zijn. Ook zijn magnetische structuren ontlekt in dit onderzoek (zie sectie 7.2), die op geen enkele andere manier waargenomen kunnen worden. Bovendien worden de "magnetische tunnels" voor de eerste maal geïnterpreteerd als voornamelijk veroorzaakt door magnetische effecten. En hoewel het afstudeerproject van Dominik Schnitzeler nog in volle gang is (sectie 7), kunnen we uit de eerste resultaten al schattingen maken over hoe sterk het grote-schaal magneetveld moet zijn.

Zo levert dit proefschrift een bijdrage aan de kennis van het interstellaire magneetveld, om uiteindelijk meer te weten te komen over het verleden (en de toekomst!) van onze zon en de Melkweg.
Curriculum vitae


Tijdens mijn promotieonderzoek heb ik deelgenomen aan workshops en conferenties in Leiden, Naramata (Canada), Bonn, Manchester, Santa Barbara, Parijs en Bologna en aan een herfstschool in Dwingeloo. Ik heb werkbezoeken afgelegd aan Dwingeloo, Heidelberg, Victoria (Canada), Penticton (Canada), Cambridge (VS) en Boulder (VS). Bovendien ben ik actief geweest in de commissie publiekscontact en de sociale commissie van de Sterrewacht Leiden en heb ik een twintigtal populaire lezingen gegeven aan niet-sterrenkundigen en scholieren.

Na mijn promotie zal ik voor drie jaar mijn werk voortzetten als postdoc op het Harvard-Smithsonian Center for Astrophysics te Cambridge (VS), in de groep van prof. B. M. Gaensler.
Nawoord

Vele mensen hebben direct of indirect tot dit boekje bijgedragen.

Van mijn familie heb ik altijd onvoorwaardelijke liefde en vertrouwen gekregen, en ik besef heel goed hoe bijzonder dat is. Ronalds steun, liefde en humor hebben veel bijgedragen aan het plezier waarmee ik op deze periode terugkijk. Schoolvrienden, studievrienden, vakantie vrienden, collega’s, klimmatten, toneelspelers en schoonfamilie hebben mij er gelukkig altijd aan herinnerd dat er meer is dan een proefschrift.

My Dutch, non-Dutch and semi-Dutch colleagues have made the Sterrewacht the warm, alive and international environment that I enjoyed so much. Furthermore, I am grateful for the excellent collaborations and for the enthusiasm and hospitality of many friends and colleagues abroad.

Zonder de geweldige ondersteuning van de computergroep en het secretariaat was alles een stuk moeilijker geweest.

Tenslotte wil ik mijn erkentelijkheid uitspreken jegens NWO, de Leidse Sterrewacht en het Leids Kerkhoven-Bosscha Fonds, voor het financieel mogelijk maken van mijn reizen en dit proefschrift.