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Galactic Magnetic Fields

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Abstract. I present a broad overview of interstellar magnetic fields in our Galaxy. Magnetic fields constitute one of the three basic constituents of the interstellar medium, the other two constituents being the ordinary matter (covered in a separate chapter of these proceedings) and cosmic rays (the main topic of the PARC Summer School). I first describe interstellar magnetic fields from an observational point of view, based on the four different methods used to detect and measure them. I then discuss, from a theoretical point of view, their origin and their impact on the interstellar medium.

Contents

1. Introduction

Numerous observations, based on starlight polarisation, Zeeman splitting, Faraday rotation, and synchrotron emission, reveal the presence of magnetic fields in our Galaxy as well as in external spirals. These magnetic fields appear to have a pressure comparable to the interstellar cosmic-ray pressure and somewhat higher than the interstellar gas pressure.

Galactic magnetic fields are undoubtedly vital for cosmic rays. These are effectively tied to magnetic field lines by the action of the Lorentz force, which sets them into gyration motion around field lines, and by the self-excitation of Alfvén waves, which limit their streaming motion along field lines to about the local Alfvén speed. At large scales, magnetic fields confine cosmic rays to the Galactic disk and control their escape into intergalactic space. At smaller scales, they maintain cosmic rays for long periods of time in the vicinity of interstellar shock waves, where they can be accelerated to high energies.

Galactic magnetic fields also have an important impact on the ordinary matter. Through the Lorentz force, they affect both its spatial distribution and its dynamics at all scales. At large scales, they help to support it against its own weight in the Galactic gravitational field. At smaller scales, they oppose the expanding gas motions driven by supernova explosions, they constrain the random motions of interstellar clouds, and they control the star formation process. In addition to their dynamical role, Galactic magnetic fields provide a heat source for the interstellar gas through magnetic reconnection.

Conversely, the interstellar gas has a critical impact on Galactic magnetic fields, which are basically frozen into it. Its weight confines them to the Galactic disk. Its motions are responsible not only for their present-day structure, but most likely also for their amplification and regeneration in the Galactic disk. In the dynamo theory, for instance, Galactic magnetic fields are amplified under the combined action of the large-scale Galactic differential rotation and small-scale cyclonic turbulent motions.

In this article, I present the current knowledge and understanding of Galactic magnetic fields. In Section 2, I review their observational status; in Section 3, I address the question of their origin; and in Section 4, I discuss their impact on the interstellar medium (ISM). Throughout the article, the Galactic radial and vertical coordinates are denoted by R and Z, respectively, and the adopted value for the Galactocentric radius of the Sun is $R_\odot = 8.5$ kpc.

2. Observational Tools and Diagnostics

2.1 Stellar Polarimetry

The presence of interstellar magnetic fields in our Galaxy was first revealed by the observational discovery of linear polarization of starlight. This polarization can be explained in terms of selective extinction by elongated dust grains that are at least partially aligned by an interstellar magnetic field. More specifically, rapidly spinning paramagnetic grains tend to make their spin axis coincide with their short axis and orient the latter along the magnetic field. Since dust grains preferentially block the component of light with polarization vector **E** parallel to their long axis, the light that passes through is linearly polarized in the direction of the magnetic field.

Thus, stellar polarimetry acquaints us with the direction of the interstellar magnetic field within a few kpc from the Sun. This method shows that the local field is parallel to the Galactic plane and that it points nearly in the azimuthal direction. Its inclination angle with respect to the azimuthal was recently evaluated at $\simeq 7.2^\circ$ (Heiles 1996), i.e., somewhat less than the standard pitch angle of the Galactic spiral pattern ($\simeq 12^\circ$).

2.2 Zeeman Splitting

The Zeeman splitting of a given atomic or molecular line occurs in the presence of an external magnetic field, whose interaction with the magnetic moment of the valence electrons causes a subdivision of certain electronic energy levels. The amplitude of the Zeeman splitting, $\Delta\nu$, is directly proportional to the magnetic field strength, B, so that, in principle, it suffices to measure $\Delta\nu$ in order to obtain B in the region of interest. Unfortunately, in the case of the widely-used 21-cm line of atomic hydrogen (denoted by H I, as opposed to H II for ionized hydrogen), $\Delta\nu$ is usually so small compared to the line width that it cannot be measured in practice. The way to circumvent this problem is to observe instead the difference between the two circularly polarized components of the 21-cm radiation, which directly yields the line-of-sight component of the magnetic field, $B_\|$.

A vast body of Zeeman-splitting observations has now built up, both for the H I 21-cm line and for several centimeter-wavelength lines of the OH molecule. ¿From a compilation of these observations, it emerges that interstellar magnetic fields have typical strengths of a few μG in regions with gas density $n \simeq 1 - 100$ cm^{-3}, and that they display only a slight tendency for B to increase with increasing n; this tendency is a little more pronounced in the higher-density range $n \simeq 10^2 - 10^4$ cm^{-3}, where field strengths may reach up to a few tens of μG (Troland & Heiles 1986; Myers *et al.* 1995; Crutcher 1999).

It is clear that Zeeman-splitting measurements are biased toward interstellar regions with high H I column densities and narrow 21-cm line widths, that is, toward cold neutral clouds.

2.3 Faraday Rotation

A linearly polarized electromagnetic wave propagating along the magnetic field of an ionized medium can be decomposed into two circularly polarized modes: a right-hand mode, whose **E** vector rotates about the magnetic field in the same sense as the free electrons gyrate around it, and a left-hand mode, whose **E** vector rotates in the opposite sense. As a result of the interaction between the **E** vector and the free electrons, the right-hand mode travels faster than the left-hand mode, and, consequently, the plane of linear polarization experiences a rotation – known as Faraday rotation – as the wave propagates. The angle by which the polarization plane rotates is equal to the wavelength squared times the rotation measure,

$$\mathrm{RM} = \mathcal{C} \int_0^L n_e \, B_\parallel \, ds \ , \tag{1}$$

where the numerical constant is given by $\mathcal{C} = 0.81$ rad m^{-2} when the free-electron density, n_e, is expressed in cm^{-3}, the line-of-sight component of the magnetic field, $B_\parallel$, is in μG, and the path length, L, is in pc. Observationally, the rotation measure of a given source can be determined by measuring the polarization position angle of the incoming radiation at two or more wavelengths.

The sources of linearly polarized waves used to carry out Faraday-rotation measurements in our Galaxy are either pulsars or extragalactic radio continuum sources. Pulsars present a double advantage in this context: first, they lie within the Galaxy and have estimable distances, and second, their rotation measure divided by their dispersion measure,

$$\mathrm{DM} = \int_0^L n_e \, ds \ ,$$

directly yields the n_e-weighted average value of $B_\parallel$ along their line of sight.

Several important conclusions can be drawn from Faraday-rotation studies (see, e.g., Simard-Normandin & Kronberg 1980; Inoue & Tabara 1981; Rand & Lyne 1994; Han & Qiao 1994; Han *et al.* 1999). First, the interstellar magnetic field has uniform (or regular) and random (or irregular) components. Near the Sun, the uniform component has a strength $\simeq 1.5\ \mu$G and the random component a strength $\sim 5\ \mu$G.

Second, the uniform field is nearly azimuthal out to several kpc from the Sun. It reverses several times along the radial direction – two or three times in the inner Galaxy ($R < R_\odot$) and probably once in the outer Galaxy. These reversals have often been interpreted as evidence that the uniform field is bisymmetic (azimuthal wavenumber $m = 1$), although it is now clear that they are also consistent with an axisymmetric ($m = 0$) field. Along the vertical, the uniform field in the inner Galaxy appears to reverse somewhere near the midplane and be roughly antisymmetric in Z, while in the outer Galaxy it would be roughly symmetric in Z.

Third, the uniform field strength increases smoothly toward the Galactic center, reaching at least 4 μG at $R = 4$ kpc. In the vertical direction, it falls off with $|Z|$ at a very uncertain rate: pulsars RMs suggest a scale height ~ 0.15 kpc, whereas extragalactic-source RMs suggest a scale height ~ 1.4 kpc. The huge discrepancy between both values could be indicative of a two-layer structure, a view supported by synchrotron-emission data (see Section 2.4).

It is important to realize that Faraday rotation measures have two intrinsic limitations. (1) They essentially probe the warm ionized medium, which presumably contains most of the interstellar free electrons but occupies only a modest fraction of the interstellar volume, so that the inferred magnetic field strengths are not necessarily representative of the ISM in general. (2) They furnish only the line-of-sight component of the Galactic magnetic field, not its total strength.

2.4 Synchrotron Radiation

The rapid spiraling motion of cosmic-ray electrons about magnetic field lines generates nonthermal radiation, termed synchrotron radiation, over a broad range of radio frequencies. The synchrotron emissivity at frequency ν due to a power-law energy spectrum of relativistic electrons, $f(E) = K_e\, E^{-\gamma}$, is given by

$$\mathcal{E}_\nu \;=\; \mathcal{F}(\gamma)\ K_e\ B^{\frac{\gamma+1}{2}}\ \nu^{-\frac{\gamma-1}{2}}\ , \tag{2}$$

where $\mathcal{F}(\gamma)$ is a known function of the electron spectral index and B is again the magnetic field strength.

Beuermann *et al.* (1985) modeled the Galactic synchrotron emissivity, based on an all-sky radio continuum map at 408 MHz. Their model possesses a spiral structure similar to that observed at optical wavelengths, and it contains two radially extended disks with comparable emissivity at midplane but with very different thickness. The midplane emissivities decrease radially outward with an exponential scale length

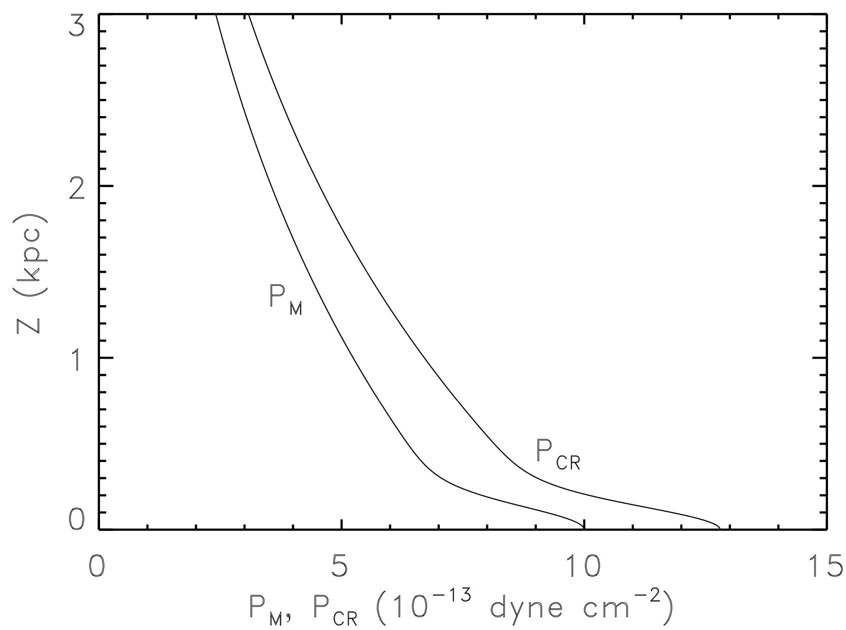

Figure 1.: *Interstellar magnetic pressure, $P_{\rm M}$, and cosmic-ray pressure, $P_{\rm CR}$, as a function of Galactic height, Z, at the solar circle ($R = R_\odot$).*

$\simeq 3$ kpc, whereas the disk thicknesses increase outward with an exponential scale length $\simeq R_\odot$. The local half-thicknesses are $\simeq 0.16$ kpc for the thin disk and $\simeq 1.5$ kpc for the thick disk.

Since the synchrotron emissivity depends on both the magnetic field strength and the spectrum of cosmic-ray electrons, it would be necessary to know one of these two quantities in order to deduce the other one from Beuermann *et al.*'s model. Unfortunately, neither quantity is well established observationally. One is thus led to follow another approach and appeal to a double assumption often made in this context, namely, the assumption that the density of cosmic-ray electrons is proportional to cosmic-ray pressure and that cosmic-ray pressure is proportional to magnetic pressure. Under these conditions, equation (2) implies that magnetic pressure, $P_{\rm M} \equiv B^2/8\pi$, and cosmic-ray pressure, $P_{\rm CR}$, are both proportional to $\mathcal{E}_\nu^{4/(\gamma+5)}$.

The last needed piece of information is the value of $P_{\rm M}$ and $P_{\rm CR}$ at a given location in the Galaxy, for instance, near the Sun. The cosmic-ray ion and electron spectra near the Sun have been determined by direct measurements – both measurements at Earth (e.g., Webber 1983a; Webber 1983b) and measurements in the heliosphere out to $\gtrsim 60$ AU by the *Voyager* and *Pioneer* spacecraft (Webber 1998) – corrected for solar modulation effects. These spectra directly lead to $(P_{\rm CR})_\odot \simeq 12.8 \times 10^{-13}$ dyne cm^{-2}. Moreover, if we introduce the parameters of the cosmic-ray electron spectrum together with the value of $(\mathcal{E}_\nu)_\odot$ at $\nu = 408$ MHz taken from Beuermann *et al.*'s model into equation (2), we find $B_\odot \simeq 5.0\ \mu$G, which translates into $(P_{\rm M})_\odot \simeq 10.0 \times 10^{-13}$ dyne cm^{-2}.

With $\gamma \simeq 2.5$, $P_{\rm M}$ and $P_{\rm CR}$ are proportional to $\mathcal{E}_\nu^{0.53}$. In particular, at the solar circle, they follow the vertical profiles displayed in Figure 1. The corresponding scale height for the magnetic field strength is $\simeq 4.5$ kpc, which is significantly greater than the value found for the uniform field component from Faraday-rotation measurements (see Section 2.3). The difference can probably be explained by the fact that the interstellar magnetic field becomes less regular away from the midplane.

3. The Origin of Galactic Magnetic Fields

3.1 Primordial Field versus Dynamo Field

Two main theories have been put forward to explain the origin of magnetic fields in our Galaxy: the primordial field theory and the dynamo theory.

In the primordial field theory, present-day Galactic magnetic fields would simply be the relics of a coherent magnetic field existing in the early Universe prior to galaxy formation. Presumably, the gas motions associated with the collapse of the protogalaxy would have compressed the lines of force of the ambient magnetic field (which are frozen into the highly conductive gas), and from then on, the differential rotation of the Galaxy would have wrapped them up about its center. In the absence of any other process, the Galactic magnetic field would be wound up 50 to 100 times at the present time, i.e., much more than indicated by observations.

The traditional way of resolving the discrepancy is to invoke magnetic diffusion (supposedly of the turbulent kind). But if magnetic diffusion is sufficiently efficient parallel to the Galactic plane to avoid a tight wind-up of magnetic field lines, it must also be sufficiently efficient perpendicular to the plane for field lines to diffuse out of the disk. Under these conditions, the Galactic magnetic field would have completely de-

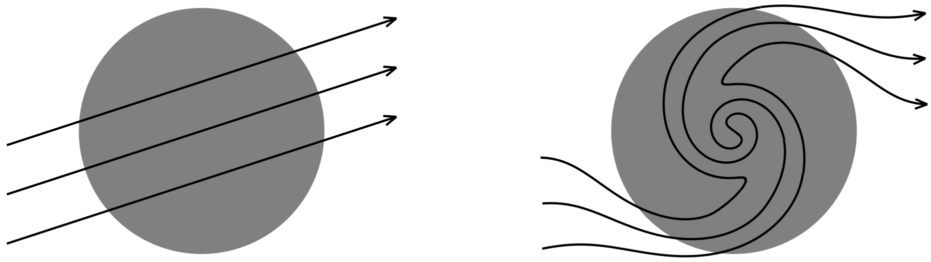

Figure 2.: *Face-on view of the Galactic disk showing the azimuthal stretching of magnetic field lines by the large-scale differential rotation.*

Figure 3.: *Oblique view of the Galactic disk illustrating the alpha-effect arising from small-scale cyclonic turbulent motions.*

cayed away by now,[1] without ever reaching the observed strengths of a few μG (Rosner & DeLuca 1989).

This short reasoning suggests that an additional mechanism partakes in the generation of Galactic magnetic fields. In the dynamo theory, this additional mechanism is due to small-scale turbulent motions which are cyclonic, i.e., which have acquired a preferred sense of rotation under the action of the Coriolis force. As these turbulent motions stretch and twist magnetic field lines, they impart to them a net rotation, whereby

[1] This conclusion could possibly become invalid in the case of ambipolar diffusion, which allows magnetic field lines, tied to the charged particles, to slip through the neutral gas.

magnetic field is created in the direction perpendicular to the prevailing field. This process has been termed the “alpha-effect”.

Thus, in the dynamo theory, the large-scale differential rotation stretches magnetic field lines in the azimuthal direction about the Galactic center (see Figure 2), while small-scale cyclonic turbulent motions regenerate, via the alpha-effect, the meridional component of the field from its azimuthal component (see Figure 3). It is the combination of these two complementary mechanisms that leads to magnetic field amplification in the Galaxy (Parker 1971; Vainshtein & Ruzmaikin 1971).

Obviously, the operation of the Galactic dynamo requires a seed magnetic field to initiate the amplification process. Several possibilities have been advanced regarding the nature of this seed field (see Rees 1987). Briefly, the seed field could have a pregalactic origin (as in the primordial field theory), or it could arise in the protogalaxy as a result of charge separation due to electrons interacting with the microwave background photons (battery effect), or else it could originate in the first generation of stars and be expelled into the interstellar medium by their winds and/or supernova explosions.

3.2 The Dynamo Equation

In the dynamo theory, the time evolution of the large-scale Galactic magnetic field is governed by the dynamo equation,

$$\frac{\partial\langle\mathbf{B}\rangle}{\partial t} = \nabla \times \Big(\langle\mathbf{v}\rangle \times \langle\mathbf{B}\rangle\Big) + \nabla \times \boldsymbol{\mathcal{E}} \,, \tag{3}$$

where $\mathbf{B}$ is the magnetic field, $\mathbf{v}$ is the velocity field,

$$\boldsymbol{\mathcal{E}} \equiv \langle\delta\mathbf{v} \times \delta\mathbf{B}\rangle \tag{4}$$

is the electromotive force due to turbulent motions, angle brackets denote large-scale (or ensemble-averaged) quantities, and the symbol δ denotes small-scale turbulent quantities (e.g., Steenbeck *et al.* 1966). The first term on the right-hand side of equation (3) represents the effect of the large-scale velocity field (including chiefly Galactic rotation) on $\langle\mathbf{B}\rangle$ and the second term represents the effect of small-scale turbulent motions.

In general, $\boldsymbol{\mathcal{E}}$ can be expressed as a linear function of $\langle\mathbf{B}\rangle$ and its first-order spatial derivatives:

$$\mathcal{E}_i = \alpha_{ij}\,\langle B_j\rangle + \beta_{ijk}\,\frac{\partial\langle B_j\rangle}{\partial x_k} \,. \tag{5}$$

The so-called alpha-tensor, α_{ij}, embodies not only the alpha-effect but also the effective advection of magnetic field by turbulent motions, whereas the tensor β_{ijk} describes turbulent magnetic diffusion (e.g., Moffatt 1978).

In the case of the Galactic disk, where the interstellar parameters and the sources of turbulence vary essentially in the vertical direction, the alpha-tensor takes on the form

$$\alpha_{ij} = \begin{pmatrix} \alpha_R & -v_{\rm esc} & 0 \\ v_{\rm esc} & \alpha_\Phi & 0 \\ 0 & 0 & \alpha_Z \end{pmatrix} \tag{6}$$

in a cylindrical reference frame $(\widehat{e}_R, \widehat{e}_\Phi, \widehat{e}_Z)$ with origin at the Galactic center and $\widehat{e}_Z$ perpendicular to the Galactic plane. The diagonal components, α_R, α_Φ, and α_Z, give the effective rotational velocity associated with the alpha-effect when $\langle \mathbf{B} \rangle$ is radial, azimuthal, and vertical, respectively, and the off-diagonal component, $v_{\rm esc}$, represents the effective vertical velocity at which $\langle \mathbf{B} \rangle$ is advected by turbulent motions (Ferrière 1993a). The diffusivity tensor, for its part, can be written as

$$\beta_{ijk} = \beta_h \, (\epsilon_{ijR} \, \delta_{kR} + \epsilon_{ij\Phi} \, \delta_{k\Phi}) + \beta_v \, \epsilon_{ijZ} \, \delta_{kZ} \ , \tag{7}$$

where β_h and β_v are the horizontal and vertical turbulent magnetic diffusivities, respectively, δ_{ij} is the unit tensor, and ϵ_{ijk} is the three-dimensional permutation tensor (Ferrière 1993b).

The rotation curve of the Galaxy is reasonably well established observationally, mainly thanks to H I and CO velocity measurements. In contrast, the dynamo tensors, α_{ij} and β_{ijk}, are poorly constrained observationally, and their determination mostly relies on theoretical models of interstellar turbulence. For example, Ferrière (1998) calculated α_{ij} and β_{ijk} in the framework of an axisymmetric model in which the turbulent motions responsible for dynamo action are driven by supernova explosions (see Figure 4).

3.3 Numerical Solutions

Several authors have set out to solve the Galactic dynamo equation (3) numerically (e.g., Donner & Brandenburg 1990; Brandenburg *et al.* 1992; Panesar & Nelson 1992; Schultz *et al.* 1994; Elstner *et al.* 1996; Ferrière & Schmitt 2000). Here we summarize the main conclusions that can be drawn from their papers.

When the parameters describing the velocity field (the large-scale velocity, $\langle \mathbf{v} \rangle$, and the dynamo tensors, α_{ij} and β_{ijk}) are prescribed, the dynamo equation is linear in $\langle \mathbf{B} \rangle$, so that the computed $\langle \mathbf{B} \rangle$ either grows or decays exponentially with time. Solutions of the linear dynamo equation are customarily classified into modes with given azimuthal symmetry (denoted by the value of the azimuthal wavenumber, m, i.e., $m = 0$ for

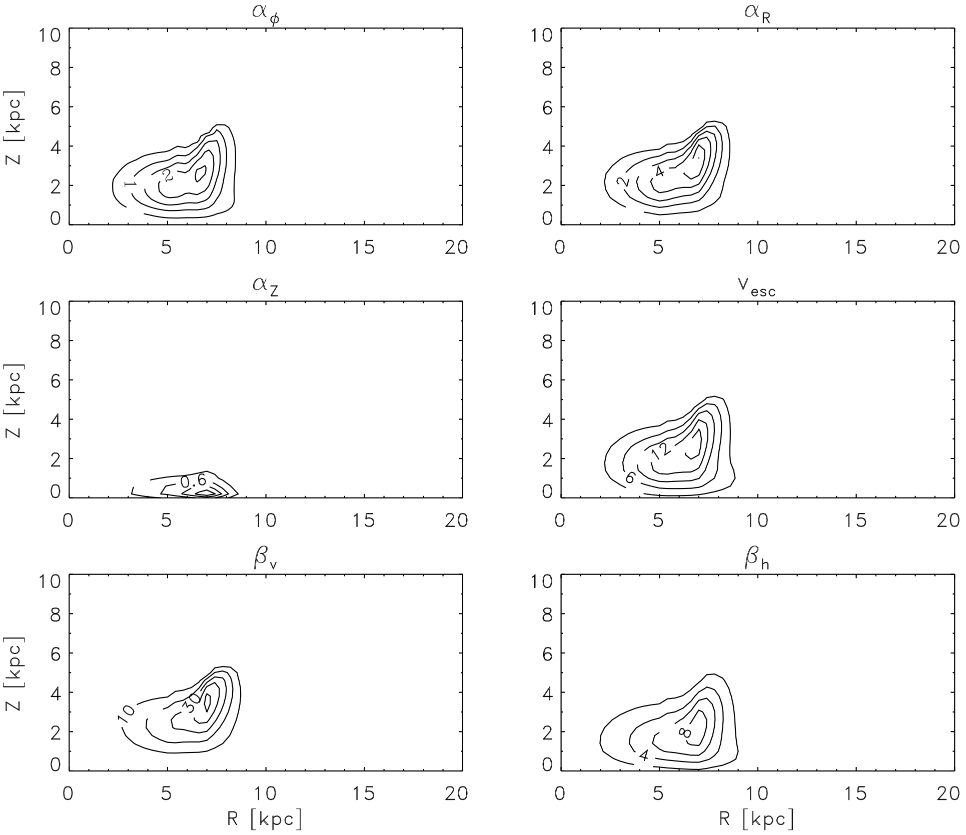

Figure 4.: *Contour lines of the four different components of the alpha-tensor (expressed in km s^{-1}) and of the two different components of the diffusivity tensor (expressed in* 10^{26} *cm^2 s^{-1}) in a given meridional plane of our Galaxy.*

axisymmetric modes, $m=1$ for bisymmetric modes ...) and vertical symmetry (indicated by the letter S or A according to whether the mode is symmetric or antisymmetric with respect to the midplane).

As far as the azimuthal symmetry is concerned, an important finding systematically emerges from all studies, namely, when the input parameters are axisymmetric, $m = 0$ modes are always easier to excite than $m=1$ modes. The physical reason is easily understood: for bisymmetric modes, the azimuthal stretching of magnetic field lines by the large-scale differential rotation has a destructive, rather than amplifying, effect; as a result, the magnetic field rapidly vanishes from the differentially-rotating parts of the Galaxy, while it undergoes a slower exponential decay in the rigidly-rotating innermost region. There exist, nevertheless, several factors that play in favor of $m=1$ modes, such as the absence of a large-scale shear, the thinness of the galactic disk, anisotropies in the alpha-tensor, and, of course, azimuthal variations in α_{ij} and β_{ijk} (arising, for instance, from the underlying spiral structure or from a gravitational encounter with a nearby galaxy).

The situation regarding the vertical symmetry is not quite as clear-cut. Let us first note that for axisymmetric magnetic fields, the azimuthal component dominates by more than one order of magnitude, as a direct consequence of the large-scale differential rotation being much more efficient at generating azimuthal field than the alpha-effect at generating meridional field. Under typical galactic conditions, the S0 mode grows faster than the A0 mode. This preference results from the characteristic disk geometry, which tends to favor quadrupolar fields (especially when the azimuthal component dominates), in contrast to the spherical geometry prevailing in the Sun, which tends to favor dipolar fields. However, under certain circumstances, the A0 mode is found to be preferred, in particular, when the galactic disk is sufficiently thick, when the alpha-effect is active and strong enough in the halo, when dynamo action is dominated by the alpha-effect near the galactic center, when differential rotation is weak, or when the alpha-tensor possesses large anisotropies.

In its standard temporal behavior, too, the Galactic magnetic field contrasts with the solar magnetic field. Indeed, whereas the latter is observed to oscillate in time with a 22-year period, the former is typically found to grow monotonously with time at a slow exponential rate $\lesssim 2\ \mathrm{Gyr}^{-1}$. Not surprisingly, certain circumstances can cause the Galactic magnetic field to become oscillatory; these include all the circumstances favoring the A0 mode (see previous paragraph) as well as a strongly reduced escape velocity and a large-scale rotation rate decreasing away from the midplane. Furthermore, the A0 mode is generally the first to turn oscillatory, and it is only when the A0 and S0 modes have both turned oscillatory that the A0 mode is liable to grow faster.

The main limitation of the linear results discussed above resides in the faulty underlying assumption that α_{ij} and β_{ijk} are time independent. In reality, the dynamo parameters evolve in the course of time, insofar as they depend on basic interstellar parameters such as the large-scale rotation velocity, the supernova rate, the ambient interstellar pressure (including a magnetic component), the interstellar mass density, and the magnetic field itself, all of which vary with time in a complicated interrelated manner. In consequence, the evolution of $\langle \mathbf{B} \rangle$ is coupled to that of the other interstellar parameters.

Although the full problem has never been solved self-consistently, there have been attempts to go beyond the purely linear approach (Jepps

1975; Brandenburg *et al.* 1992). These attempts generally consist of allowing the components of α_{ij} (and sometimes also β_{ijk}) to depend on $\langle \mathbf{B} \rangle$ via a multiplicative quenching function, $q(\langle \mathbf{B} \rangle)$, meant to represent the back-reaction of the Lorentz force on the turbulent motions responsible for the dynamo process. The quenching function is generally assigned a convenient form, which has the expected property of decreasing to zero as the magnetic field strength increases to infinity. A common choice is

$$q(\langle \mathbf{B} \rangle) = \frac{1}{1 + \left(\frac{|\langle \mathbf{B} \rangle|}{B_{\rm eq}} \right)^n} \;, \tag{8}$$

where $B_{\rm eq}$ is the equipartition field strength (such that $B_{\rm eq}^2/8\pi$ equals the interstellar gas pressure) and n is usually set to 2.

Evidently, nonlinear solutions differ from their linear counterparts in several respects. First, magnetic field growth is exponential only at early stages (provided there is indeed positive growth); it begins to saturate when $|\langle \mathbf{B} \rangle|$ approaches $B_{\rm eq}$, and the fully saturated field strength generally does not exceed $\sim 2\, B_{\rm eq}$. The time to saturation, which depends on the adopted seed field strength, is typically on the order of a few tens of Gyr. Second, since quenching and the ensuing reduction in field amplification are, by construction, more severe in stronger-field regions, the final magnetic configuration tends to be more smoothed out (i.e., less peaked) than in linear models. Third, nonlinear interactions between different magnetic modes make it possible to amplify certain modes that would otherwise decay away. These interactions can maintain mixed-parity configurations for extended periods of time, and in some cases, the magnetic field eventually evolves toward a state whose parity differs from the preferred linear parity.

4. The Impact of Galactic Magnetic Fields on the ISM

4.1 Dynamic Role

The Galactic magnetic field acts on the interstellar matter through the Lorentz force. Of course, the field acts directly on the charged particles only, but its effect is then transmitted to the neutrals by ion-neutral collisions. Apart from the densest parts of molecular clouds, whose ionization degree is exceedingly low, virtually all interstellar regions are sufficiently ionized for their neutral component to remain tightly coupled to the charged component and, hence, to the magnetic field.

At large scales, the Galactic magnetic field helps to support the ordinary matter against its own weight in the Galactic gravitational potential, and it confines cosmic rays to the Galactic disk. In this manner, both magnetic fields and cosmic rays partake in the overall hydrostatic balance of the ISM and influence its stability. Boulares & Cox (1990) were the first authors to fully appreciate the importance of magnetic fields and cosmic rays in the hydrostatic balance. By the same token, they managed to solve the long-standing problem of apparent mismatch between the total interstellar pressure at a given point and the integrated weight of overlying interstellar material: by adopting higher magnetic and cosmic-ray pressures than previously estimated, they were able to bring the total pressure at low $|Z|$ into agreement with the integrated weight, and by including the magnetic tension force at high $|Z|$, they could explain why the weight integral falls off faster than the total pressure.

As magnetic and cosmic-ray pressures inflate the gaseous disk, they tend to make it unstable to a generalized Rayleigh-Taylor instability, now known in the astrophysical community as the Parker instability (Parker 1966). When this instability develops, magnetic field lines ripple, and the interstellar matter slides down along them toward the magnetic troughs, where it accumulates. This whole process, it has been suggested, could give birth to new molecular-cloud complexes and ultimately trigger star formation (Mouschovias *et al.* 1974; Elmegreen 1982).

At smaller scales, the Galactic magnetic field affects all kinds of turbulent motions in the ISM. Of special importance is its impact on supernova remnants and superbubbles (see Tomisaka 1990; Ferrière *et al.* 1991; Slavin & Cox 1992). First, the background magnetic pressure acting on their surrounding shells directly opposes their expansion. Second, the magnetic tension existing in the field lines swept into the shells gives rise to an inward restoring force, while the associated magnetic pressure prevents the shells from fully collapsing and, therefore, keeps them relatively thick. Third, the enhanced external "signal speed" causes the shells to merge earlier than they would in an unmagnetized medium. All three effects conspire to lower the filling factor of hot interstellar gas.

The Galactic magnetic field also constrains the random motions of interstellar clouds. The latter are magnetically connected to their environment, namely, to the intercloud medium and possibly to neighboring clouds, through the magnetic field lines that thread them. When a given cloud moves relative to its environment, these field lines get deformed, and the resulting magnetic tension force modifies the cloud's motion, transferring part of its momentum to its environment (Elmegreen 1981). Likewise, angular momentum can be transferred from a rotating cloud to its environment by magnetic torques. This mechanism is particularly relevant to the star formation process, as it allows the contracting protostellar cores to get rid of angular momentum (Mouschovias & Morton 1985).

Finally, the Galactic magnetic field plays a crucial role in the support of molecular clouds against their self-gravity and in the eventual gravitational collapse of protostellar cores (Shu *et al.* 1987). The magnetic support of molecular clouds is essentially provided by magnetic pressure gradients in the directions perpendicular to the mean field and, presumably, by nonlinear Alfvén waves in the parallel direction. In the case of protostellar cores, magnetic support is insufficient to prevent their ultimate gravitational collapse. This is generally because their ionization degree is so low that neutrals are not perfectly tied to magnetic field lines, which enables them to drift inwards under the pull of their self-gravity and eventually form stars.

4.2 Energetic Role

In most regions of interstellar space, the electrical conductivity, σ, is extremely high, with the twofold consequence that magnetic field lines are frozen into the plasma and that ohmic dissipation is completely negligible. There exist, however, localized regions where magnetic field gradients are very steep, so that the electric current density, $\mathbf{J} = \frac{c}{4\pi} \nabla \times \mathbf{B}$, is very strong, and where the effective electrical conductivity is greatly reduced with respect to its classical Coulomb value (for instance, an anomalous conductivity may result from microscopic wave fluctuations which increase the effective collision frequency through wave-particle interactions). The combination of strong $\mathbf{J}$ and reduced σ entails both a partial loss of the magnetic field frozen-in property and a resistive dissipation of magnetic energy.

In particular, when field lines of opposite polarities are forced to approach one another, a thin transition layer develops, in which the magnetic field gradient steepens and, accordingly, the current density rises, to the point where field lines decouple from the plasma, reconnect with field lines of the opposite polarity, and leave the transition layer sideways. The reconnection zone is subject to intense ohmic dissipation, with the dissipated magnetic energy being essentially converted into plasma thermal energy.

Thus, magnetic reconnection constitutes a source of heating – and, at the same time, ionization – for the ISM. This heating/ionization source could even be dominant in the Galactic halo, which is not easily accessible to the powerful radiation from the luminous O and B stars. Based on this idea, Birk *et al.* (1998) invoked magnetic reconnection within thin current filaments assumed present in the halo to solve the ionization/heating problem of the warm ionized medium at high $|Z|$. It was also suggested by Zimmer *et al.* (1997) that interaction regions between high-velocity clouds traveling through the halo and the surrounding halo gas could be the sites of intense magnetic reconnection and, hence, plasma heating.

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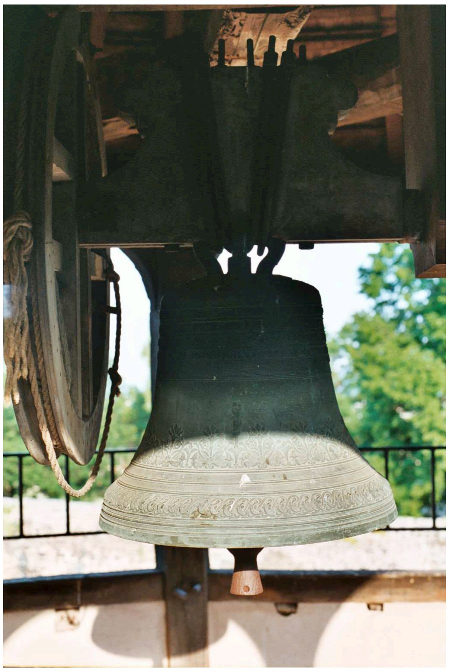

Cloche du château de Goutelas

Entrée de la chapelle

Fresque restaurée à l'intérieur de la chapelle