# MAGNETIC FIELDS ON SOLAR–LIKE STARS: THE FIRST DECADE

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ABSTRACT. I review the progress made over the past decade in the measurement of magnetic fields on solar–like stars. I describe the evolution of magnetic analysis techniques, summarize our current understanding of stellar magnetic properties, and outline some future research directions.

## 1. Introduction

It has now been nearly a decade since Robinson *et al.* (1980) opened a new era of stellar research by making the first measurements of magnetic fields on solar–like stars. In light of this anniversary it is appropriate to review the field's development and its current analysis techniques, summarize the present state of knowledge in this field, and look ahead to future developments. I also encourage the reader to seek out recent reviews by Saar (1987b), Gray (1988), and Linsky (1989).

The reasons for studying magnetic fields on solar–like stars rise directly from the study of magnetic fields on the Sun. The entire family of atmospheric features on the Sun (plages, spots, granules, prominences, flares, etc.) is related to, or significantly affected by, the presence of magnetic flux. The inferred presence of similar features on cool stars suggests comparable, magnetically–related origins (e.g., Linsky 1985). The million degree corona of the Sun requires magnetic fields for both heating (by direct and indirect means) and confinement of the hot plasma. Observations of stellar coronae, with emission often orders of magnitude larger than the Sun, reinforce the concept that magnetic fields are ubiquitous in cool stars (e.g., Golub 1983). These are but two examples. Clearly, data on magnetic fields is vital to understand fully the physical structure, energy balance and evolution of the stellar atmospheres which the fields permeate.

## 2. Evolution of Analysis Methods

Unfortunately, the solar analogy breaks down when one seeks to employ solar *techniques* for measuring magnetic fields to the stars. Lack of spatial resolution, combined with the (likely) dipolar structure of the magnetic regions, are the the fundamental problems. They conspire to almost completely cancel stellar circular polarization (e.g., Borra *et al.* 1984), and reduce



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the linear polarized signal from cool stars to a few parts in $10^3$ or less, typically (Huovelin *et al.* 1988). Magnetic broadening of the unpolarized line profiles remains detectable, but the splitting ($\Delta v_B$) is, in general, considerably less than the intrinsic line width ($\Delta v$) at optical wavelengths ($\Delta v \approx 4$ km s$^{-1}$ at 600 nm in the Sun, compared with $\Delta v_B = 1.4\times10^{-7} g_{eff} \lambda B = 2$ km s$^{-1}$ for $g_{eff} = 2.5$ and B = 1 kG). Furthermore, the Zeeman broadening effect is diluted by the (usually dominant) contributions from non–magnetic regions on the stellar surface. Finally, the lines chosen must be free of blends, and non–magnetic broadening in the lines should be low. As a result of these requirements, data must have high spectral resolution ($\Delta\lambda \leq 2\Delta\lambda_B$), high S/N (> 100–200), and the star should have $\Delta v_B/\Delta v \geq 0.2$ (e.g., $v \sin i \leq 10$ km s$^{-1}$ for a G dwarf at 600 nm). Unfortunately, this last constraint does not allow measurement of some of the most interesting, rapidly rotating stars! Some of the above difficulties can be circumvented by observing in the infrared (pioneered by Giampapa *et al.* 1983) to take advantage of the $\lambda$ dependence of $\Delta v_B$. Ultimately, however, the $v \sin i$ constraint is an erent limitation of the method.

Despite these constraints, by the mid–1970's data of the necessary quality could readily be obtained, thanks to advances in electronic detector technology. New analysis methods were now needed. The breakthrough came when Robinson *et al.* (1980), using the Fourier ratio technique pioneered in solar work by Tarbell and Title (1977), discovered a solar–like field covering a substantial fraction of the active G dwarf, $\xi$ Boo A. Marcy (1982) soon followed with a similar analysis in the wavelength domain. These methods, and all those developed to date, assume that an observed line profile, $F_{obs}$, can be interpreted with a two–component model, $F_{obs} = fF_m(B) + (1\text{-}f)F_q(B=0)$, where $F_m$ is the profile arising from magnetic regions with a mean field strength B covering a fraction f of the surface, and $F_q$ represents the profile in the field–free (B=0) regions. Methods differ primarily on how $F_m$ and $F_q$ are computed (or derived) and how the resulting magnetic parameters f and B are obtained.

The first analysis methods assumed $F_m$ could be modeled as the appropriately weighted sum (Marcy 1982) or convolution (Robinson 1980) of a triplet of optically thin lines split by a field strength B. The weights reflected the Sears relations for a "disk–averaged" angle ($\gamma$) between the magnetic field and the line–of–sight (e.g., Marcy 1982, 1984 used $\overline{\gamma} = 34°$). Lines with low $g_{eff}$ from the same star were used to simulate these magnetic components and $F_q$. Gray (1984) retained these basic assumptions but extended the Fourier ratio technique to analyze several lines simultaneously by first removing the underlying unsplit profile (= $F_q$) of each line through a radiative transfer calculation.

Some problems began to appear, however. The filling factors on many stars seemed unrealistically large (nearly 90 % for the moderately active K dwarf $\epsilon$ Eri, for example). Large magnetic fluxes were recorded for stars with widely different rotation rates and activity levels. Indeed, Gray (1985) showed that the detections made up to that time showed a constant magnetic flux density (i.e., the product, fB) for all stars, with the single exception of the Sun. Taken at face value, this result seemed to suggest two possibilities: Either theories regarding the generation of stellar magnetic fields and their roles in heating chromospheres and coronae were wrong, or something was amiss with magnetic measurements.

It now appears that shortcomings in the analysis played an important role in these problems (Hartmann 1987; Saar 1988a; Basri and Marcy 1988). None of the analyses to that time included treatment of radiative transfer effects *in the Zeeman components*

*themselves*; the construction of $F_m$ was made in the optically thin, weak–line limit. This assumption required large filling factors to duplicate observed "saturation" in optically thick line cores, especially in K dwarfs where the lines employed were generally stronger. When combined with some subtle selection effects (faster rotating, more chromospherically active G dwarfs and slower rotating, less active K dwarfs made up most of the sample; see Stepień 1987), these problems led the early analyses to infer fB = constant.

Clearly, a refined analysis including a radiative transfer model for $F_m$ was needed. Steps in this direction were already taken by Marcy and Bruning (1984), who developed a radiative transfer model to bootstrap comparison between low and high $g_{eff}$ lines of substantially different excitation potential. Full integration of magnetic line transfer effects into the analysis was accomplished by Saar and coworkers (Saar and Linsky 1985; Saar *et al.* 1986a; Saar 1988a), who developed methods which included simple magnetic radiative transfer effects (Unno 1956), the full Zeeman patterns, and some compensation for line blends. These models, however, used convolutions (Gray 1976) to describe velocity broadening, and still required the assumption of an average $\overline{\gamma}$. Bruning (1984) noted that disk integrations are preferred over convolutions for computation of rotational broadening and Saar (1988b) showed use of convolutions could produce errors in f and B values (see also Landolfi *et al.* 1989). The physics of the line transfer was also quite simple, employing a Milne–Eddington atmosphere with a linear source function and all other variables independent of depth.

These shortcomings have also been addressed recently. Basri and Marcy (1988) and Marcy and Basri (1989) have further improved the analysis by numerically integrating the Unno (1956) equations in realistic model atmospheres (rather than using the Milne–Eddington approximation). The latest models also employ full disk integration of intensity profiles to obtain the flux (Saar *et al.* 1989; Marcy and Basri 1989). This step simultaneously eliminates the need to assume an $\overline{\gamma}$, since the angle is automatically accounted for in the disk integration. Known blends can be treated simultaneously by direct line synthesis.

Mathys and Solanki (1989), however, have taken a very different approach. Their technique, based on the solar multi–line regression analysis of Stenflo and Lindegren (1977), correlates line areas measured below the half depth point and the line width at this level with parameters such as excitation potential and a Zeeman broadening term. The Zeeman term, proportional to $fB^2$, can then be separated into f and B by comparison with similar studies of the areas and widths for different line depths. Typically, large numbers of lines ($\approx 40$) are used in the analysis. By sorting lines of varying excitation potential and Zeeman pattern, some crude information on the temperature and average orientation ($\overline{\gamma}$), respectively, of the magnetic regions can potentially be determined. The method is also considerably simpler than detailed model calculations for all the lines. The functional form assumed for the regression equation, however, is physically somewhat unclear, and the choice of its terms is not straightforward. Thus, while the technique appears quite promising, additional calibration and tests are needed.

## 3. A Summary of Current Knowledge

Using the above techniques, our understanding of stellar magnetic properties has grown steadily over the last decade. In this section I summarize the current knowledge of stellar field strengths and filling factors, and present a fresh analysis of the most recent results.

First, Zeeman broadening is definitely present in measurable quantities on many G and K dwarfs (e.g., Robinson *et al.* 1980; Marcy 1984; Gray 1984), often at levels far exceeding that seen in the Sun. It has not yet been observed in F dwarfs (Gray 1984) or in any bright giants or supergiants. It is also usually not seen in subgiants and normal giants (Marcy and Bruning 1984; Gray and Nagar 1985) unless the stars are active RS CVn variables (e.g., Giampapa *et al.* 1983). And while the presence of Zeeman broadening is small, often debatable, and always difficult to interpret in optical spectra, it is undeniably present in infrared spectra of M dwarfs, where full splitting patterns are visible (Saar and Linsky 1985; Saar *et al.* 1987). There is also clear evidence for range in Zeeman broadening at each spectral type, since there are clear *non*–detections for G (Marcy 1984; Gray 1984; Mathys and Solanki 1989), K (Saar 1987a; Saar *et al.* 1987; Marcy and Basri 1989) and M (Saar *et al.* 1987) dwarfs as well. Strong magnetic fields have been detected in stars as cool as M4.5 (EV Lac; Saar *et al.* 1987), indicating that dynamo generation of magnetic flux is still effective even in stars which are almost fully convective.

To explore trends in magnetic parameters f and B in more detail, I have compiled a critically selected group of recent magnetic measurements in Table 1. The list includes virtually all measurements to date which have been derived using radiative transfer methods. This restriction makes the data set more homogeneous, and also excludes probable systematic effects in the earlier measurements (see section 2). In addition, a few optical M dwarf measurements (Bruning *et al.* 1987) were excluded due to probable blend problems (e.g., Figure 1 of Saar 1988b), and a few results based on lower S/N optical data (in Saar 1987) were also excluded. I included the infrared detections of $\lambda$ And (Giampapa *et al.* 1983; Gondoin *et al.* 1986) and the results of Mathys and Solanki (1989). In the latter case, B was computed from the mean of the range given (their table 4), and then I assumed $f \equiv \bar{f} = (\sqrt{f}B/\overline{B})^2$. Rotational periods were gathered from Noyes *et al.* (1984) and Pettersen (1989), and relations in Stepień (1989) and Basri (1987) were used to compute the convective turnover time, $\tau_c$. Two measurements are given for a single star in several cases where different investigators disagreed on the magnetic parameters (e.g., HD 131156A = $\xi$ Boo A). When this occurs, both are plotted and connected with a dotted line.

One of the first things apparent is a general tendency for magnetic field strengths to increase with increasing B–V (i.e., towards later spectral types). This trend, combined with the lack of detections for F dwarfs and most giants and subgiants (likely due to low B; Marcy and Bruning 1984), and the low field strengths seen on the active giant, $\lambda$ And, strongly suggests a relationship between B and some intrinsic property of the stellar atmosphere. Saar and Linsky (1986a) found a tight correlation between B and the equipartition "pressure–balancing" magnetic field, $B_{eq} \propto P_{gas}^{0.5}$, where $P_{gas}$ is the photospheric gas pressure. I estimated $P_{gas}$ as a function of $T_{eff}$ and gravity by using atmospheres of Kurucz (1979 and unpublished), supplemented with models derived by Mould (1976) for M dwarfs. The gas pressure was determined at optical depth, $\tau_{5000} = 1$, where $P_{gas}(\odot)$ yields $B_{eq} \approx 1.5$ kG (note that the exact choice of $\tau_{5000}$ is not critical, since we use $P_{gas}$ only to scale relative

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Table 1: A Critical Selection of Recent Magnetic Field Determinations

| Star ID | Sp. Type | B–V $(R-I)$ | B (kG) | f (%) | Ref. | adopted $B_{eq}$ | $P_{rot}$ (days) | $\tau_c$ (days) |
|---|---|---|---|---|---|---|---|---|
| HD 39587 | G0V | 0.59 | 1.0 | 60 | 1 | 1.4 | 5.2 | 10.0 |
| HD 190406 | G1V | 0.61 | 1.8 | 10 | 1 | 1.4 | 13.5 | 11.0 |
| HD 1835 | G2V | 0.66 | 1.4 | 32 | 1 | 1.5 | 7.7 | 13.7 |
| HD 28099 | G6V | 0.66 | 1.7 | 30 | 2 | 1.5 | 8.7 | 13.7 |
| HD 20630 | G5V | 0.68 | 1.5 | 35 | 1 | 1.6 | 9.4 | 14.7 |
| HD 10700 | G8V | 0.72 | ($\sqrt{f}$B | $\leq$0.2) | 12 | 1.7 | 31.9 | 16.9 |
| HD 131156 A | G8V | 0.76 | 1.6 | 22 | 3 | 1.7 | 6.2 | 19.0 |
| | G8V | 0.76 | 1.8 | 35 | 1 | 1.7 | 6.2 | 19.0 |
| HD 152391 | G8V | 0.76 | 1.7 | 18 | 1 | 1.7 | 11.1 | 19.0 |
| HD 3651 | K0V | 0.85 | ... | ... | 3 | 1.9 | 48 | 20.5 |
| HD 10476 | K1V | 0.84 | 1.0 | 17 | 3 | 2.0 | 38 | 20.5 |
| HD 165341 | K1V | 0.86 | 1.2 | 18 | 3 | 2.0 | 19.7 | 20.5 |
| HD 155885 | K1V | 0.86 | 1.5 | 13 | 3 | 2.0 | 22.9 | 20.5 |
| HD 22049 | K2V | 0.88 | 1.0 | 30 | 3 | 2.2 | 11.3 | 20.5 |
| | K2V | 0.88 | 1.9 | 12 | 4 | 2.2 | 11.3 | 20.5 |
| HD 115404 | K2V | 0.93 | 2.1 | 20 | 1 | 2.2 | 18.8 | 20.5 |
| HD 45088 A | K3Ve | 0.96 | 2.4 | 50 | 2 | 2.4 | 7.4 | 20.5 |
| HD 209100 | K4-5V | 1.09 | 2.6 | 13 | 12 | 2.5 | ... | 20.5 |
| HD 225732 | K6V | 1.04 | 1.8 | 20 | 7 | 2.5 | ... | 20.5 |
| HD 131156 B | K4V | 1.10 | (if 2.6 | $\leq$20) | 1 | 2.6 | 11.5 | 20.5 |
| EQ Vir | K5Ve | 1.18 | 2.5 | 80 | 5 | 2.7 | 3.9 | 20.5 |
| HD 201091 | K5V | 1.18 | (if 1.5 | $\leq$5) | 1,6 | 2.7 | 37.9 | 20.5 |
| | K5V | 1.18 | 1.2 | 24 | 3 | 2.7 | 37.9 | 20.5 |
| BY Dra | K5Ve+ | 1.19 | 2.8 | 60 | 8 | 2.7 | 3.8 | 20.5 |
| HD 201092 | K7V | 1.38 | (if 1.5 | $\leq$10) | 8 | 2.9 | 48.0 | 20.5 |
| HD 97101 | K9V | 1.35 | 1.8 | 25 | 7 | 2.9 | ... | 20.5 |
| HD 88230 | K7V | 1.37 | (if 1.5 | $\leq$10) | 8 | 2.9 | ... | 20.5 |
| | K7V | 1.37 | 0.8 | 55 | 7 | 2.9 | ... | 20.5 |
| GL 205 | M1.5V | (0.85) | (if 1.5 | $\leq$15) | 8 | 3.4 | ... | 20.5 |
| AU Mic | M1.5Ve | (0.84) | 4.0 | 90 | 8 | 3.4 | 4.8 | 20.5 |
| =AD Leo | M3.5Ve | (1.12) | 3.8 | 73 | 6,8 | 3.8 | 2.7 | 20.5 |
| GL 273 | M4V | (1.15) | (if 1.5 | $\leq$25) | 8 | 4.0 | ... | 20.5 |
| EV Lac | M4.5Ve | (1.15) | 5.2 | 90 | 8 | 4.4 | 4.4 | 20.5 |
| HD 222107 | G8III–IV | 1.01 | 0.6 | 20 | 9 | 0.6 | 20.5 | 87.3 |
| | G8III–IV | 1.01 | 1.2 | 48 | 10 | 0.6 | 20.5 | 87.3 |
| HD 17433 | K1-2IVe | 0.96 | 2.0 | 60 | 11 | 1.3 | 12.1 | 89.6 |

References: [1]Saar (1987); [2]Saar and Linsky (1986); [3]Basri and Marcy (1989); [4]Saar *et al.* (1986b); [5]Saar *et al.* (1986a); [6]Saar and Linsky (1985); [7]Bruning *et al.* (1987); [8]Saar *et al.* (1987); [9]Gondoin *et al.* (1986); [10]Giampapa *et al.* (1983); [11]Bopp *et al.* (1989); [12]Mathys and Solanki (1989)

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to the Sun). The resulting $B_{eq}$ was then determined using $B_{eq} = (P_{gas}/P_{gas}(\odot))^{0.5} B(\odot)$. The results (with $B_\odot \equiv 1.5$ kG; Harvey and Hall 1975) are given in Table 1; for G and K stars they are very similar to $B_{eq}$ derived by Zwaan and Cram (1989).

A plot of B and $B_{eq}$ (Figure 1) shows most stars clustering near or below the $B = B_{eq}$ line. I interpret this as evidence that magnetic field strengths in stellar photospheres are largely determined by a horizontal pressure balance in which gas pressure dominates, i.e., $B \leq B_{eq} \propto P_{gas}^{0.5}$. Several theories of magnetic field concentration in stellar photospheres predict this type of relation (e.g., Parker 1978; Spruit and Zweibel 1979; Galloway and Weiss 1981). One RS CVn data point lies considerably above the $B = B_{eq}$ line (Giampapa *et al.* 1983); Giampapa (1984), however, postulates that the stellar surface had an admixture of umbrae (with much higher B) in the field of view during the observation. In the case of the second RS CVn (HD 17433), which also shows a rather large field strength, enhanced turbulence may play a larger role in the pressure balance (Bopp *et al.* 1989).

Hartmann's (1987) suggestion that the $B \propto B_{eq}$ relation might be due to line opacity effects is probably not correct, since the analyses used to derive the data in Table 1 explicitly take opacity broadening into account. His comment that stars with low B may be missed due to difficulties in detection (e.g., Marcy 1982; Saar 1988a), however, is certainly valid. An equally important problem is the following: How valid it is to mix results of data

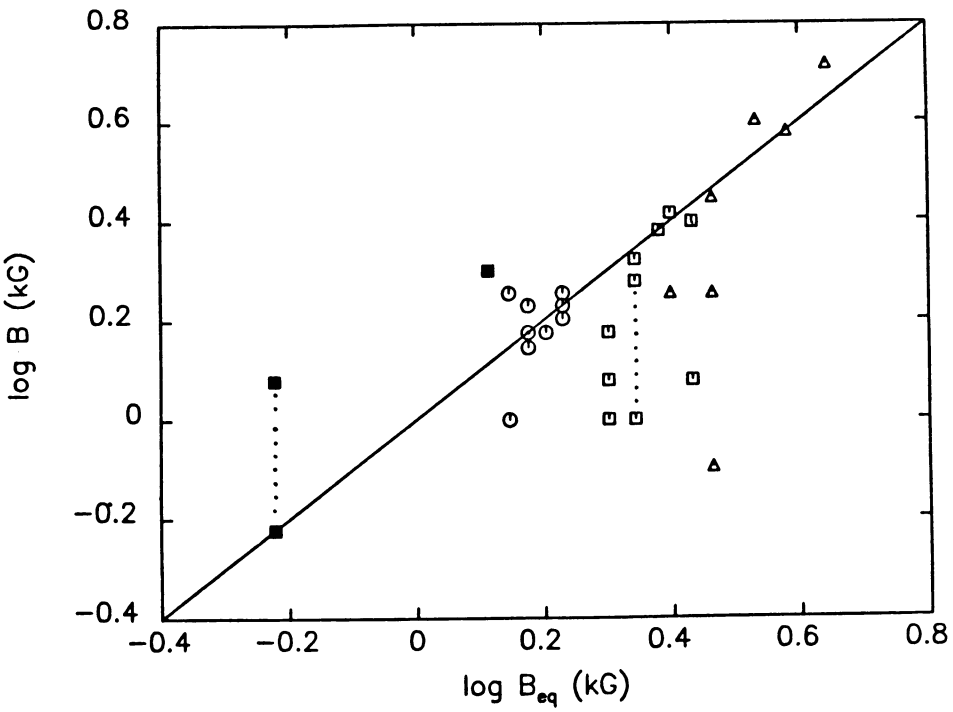


**Figure 1.** B versus $B_{eq} \propto P_{gas}^{0.5}$. Circles, squares, triangles, and filled squares represent G, K, M dwarfs and RS CVn variables, respectively. The Sun is indicated by $\odot$. The relation $B = B_{eq}$ is shown as a solid line. $B \leq B_{eq}$ is inferred from the data.

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sets measured using different lines and wavlengths? Table 1 contains magnetic parameters derived from medium strength Ti I lines at 2200 nm (e.g., Saar and Linsky 1985), Fe I lines at 617 nm (e.g., Saar *et al.* 1986a) and 1600 nm (Gondoin *et al.* 1986) and strong Fe I lines at 846 nm (Marcy and Basri 1989). Continuum opacity (primarily $H^-$) is similar at 617 and 2200 nm, but is stronger at 846 nm and substantially weaker at 1600 nm (I thank S. Solanki for pointing this out). If all other variables are held constant, fields should appear stronger (with a lower f) at 1600 nm, where the observer sees deep into the atmosphere, and weaker (with higher f) at 846 nm, where the observer sees deep into the atmosphere. Indeed, the B values derived for K dwarfs using the 846 nm line (Marcy and Basri (1989) *are* somewhat smaller (and the f values larger) than those of derived using 617 nm. The considerably greater strength of the 846 nm feature will make some parts of its profile more sensitive to higher atmospheric levels (with lower B and higher f values) than 617 nm. Similarly, optical determinations of B differ from those at 1600 nm by 300–400 G in the Sun (e.g., Stenflo 1989). Thus, it is possible that some of the variation in measured field strengths in K dwarfs is due to the effects of differing line formation heights. Of course, differences in the analysis methods used will also affect the results. A more detailed analysis, using flux tube models (e.g., Steiner *et al.* 1986) to determine the contribution functions (e.g., Grossmann–Doerth *et al.* 1988) for magnetic lines at different wavelengths, will probably be needed to solve this puzzle.

A search for other correlations with B yields little. Plots of B versus f show no correlation (see also Saar and Linsky 1986a), verifying that fB is not a constant. Also, $\sqrt{f}B$ is not constant, as would be expected if errors in the separation of f and B were dominant (Gray 1984; Saar 1988a). No clear relation exists between B and rotation, either. Plots of B with the inverse Rossby number, $\tau_c\Omega$, are basically flat. There appears to be a weak relationship between B and $\Omega$ ($B \propto \Omega^{0.3\pm0.2}$; see also Saar 1987a), but this probably represents a selection effect (only stars with large B can be detected at large $\Omega$; Saar 1988a).

The lack of a B–rotation correlation is at first suprising; many simple dynamo models predict $B \propto \Omega$ or $\tau_c\Omega$ (e.g., Stix 1972). However, when it is understood that dynamo theories actually predict the magnetic *flux* generated, rather than field *strength*, the paradox is resolved. Marcy (1984) and Gray (1985)noted the first indications that magnetic flux measurements and stellar rotation were related. Using mostly older data, Stepień (1987) found magnetic flux correlated with $(\tau_c\Omega)^{0.5}$. In the present data set, magnetic flux density (fB) exhibits strong correlations with the stellar rotation: $fB \propto \Omega^{1.3}$ and $fB \propto (\tau_c\Omega)^{1.2}$ (Figure 2; see also Saar 1987a,b). These relations are consistent with the $fB \propto T_{eff}{}^{-n}\Omega$ relation correlation found by Marcy and Basri (1989), and are broadly consistent with many simple theoretical expectations (e.g., Schatzman 1962). The Durney and Robinson (1982) dynamo model, however, which predicts $f \propto \Omega^{2.5}$, appears to be ruled out.

Since B is not strongly correlated with rotation, the filling factor must be the cause for the fB – rotation relations. A plot of f versus $\tau_c\Omega$ confirms this (Figure 3). A least–squares fit to the Table 1 data set yields $f \propto (\tau_c\Omega)^{0.9}$ (see Linsky and Saar 1987; Saar 1987b; Saar *et al.* 1987; Stepień 1988a). The filling factor must be less than one by definition, so it is inevitable that this simple power law breaks down at some point and f "saturates". Based on the derived power law, stars beyond $\tau_c\Omega \approx 0.8$ may be in this state (see also Saar and Linsky 1986a; Linsky and Saar 1987). Thus, a dynamo model including a saturated state (e.g., Skumanich and MacGregor 1986) fits the data best. Saturated states in chromospheric

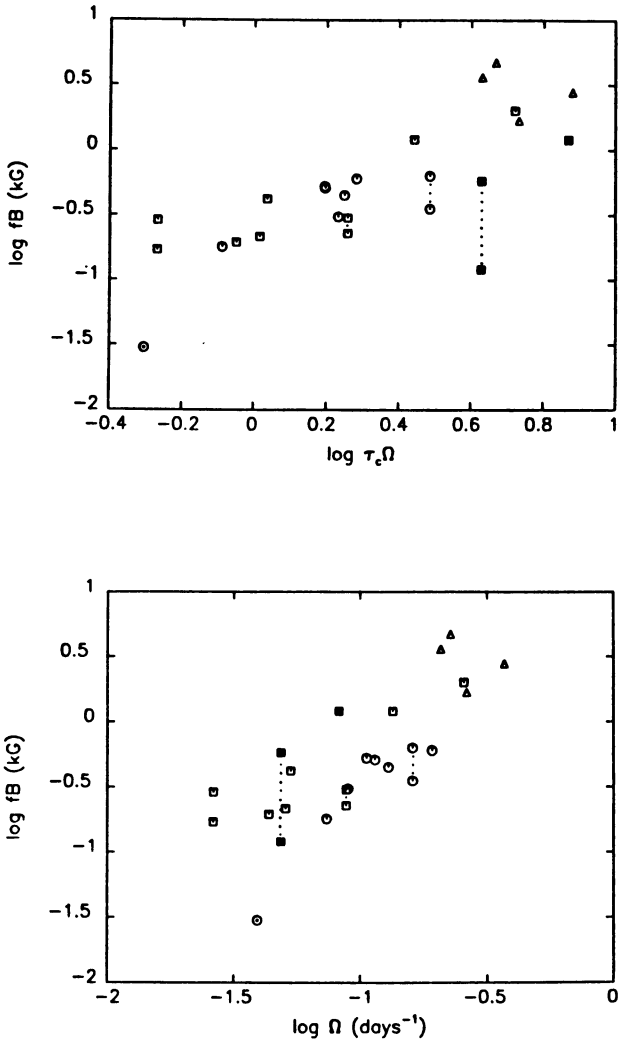


**Figure 2.** Magnetic flux density (fB) versus rotational parameters. Symbols are same as in Fig. 1. Magnetic flux density correlates with $\tau_c\Omega$ (fB $\propto (\tau_c\Omega)^{1.2\pm0.1}$; top) and $\Omega$ (fB $\propto \Omega^{1.3\pm0.1}$; bottom) largely the result of correlations between f and rotation (see Fig. 3).

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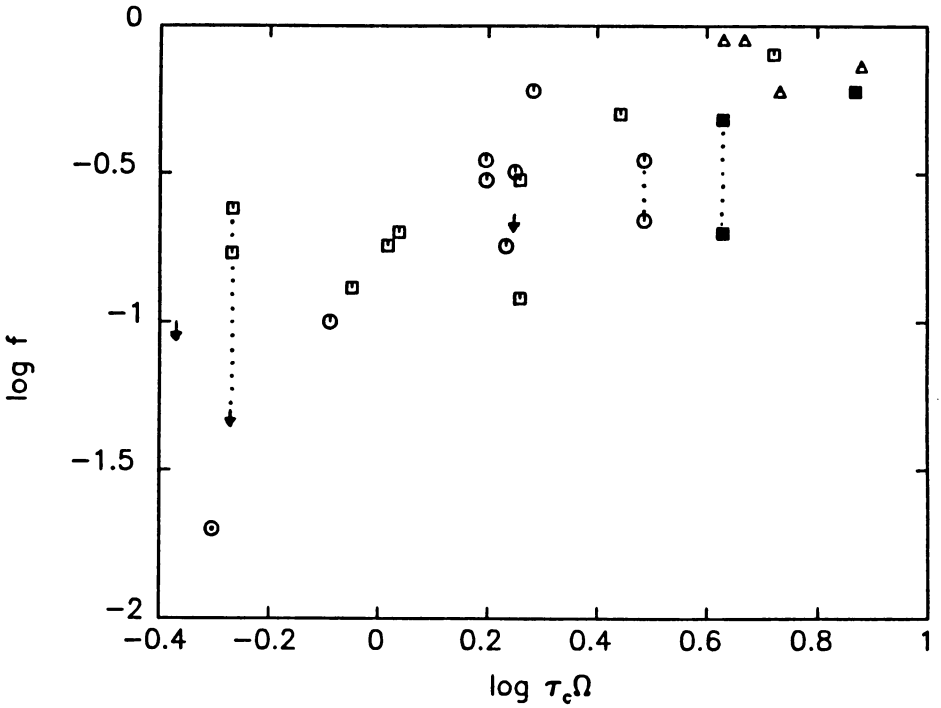


**Figure 3.** Magnetic filling factor f versus $\tau_c\Omega$. Symbols are as in Fig. 1, with upper limits to f indicated by arrows. A least-squares fit yields $f \propto (\tau_c\Omega)^{0.9\pm0.2}$. The magnetic filling factor appears to be the dominant magnetic parameter controlling rotation–activity and rotation–age relations.

and coronal emission have been recognized for some time (e.g., Vilhu 1984).

As one might anticipate based on the fB – rotation correlations, there are also indications (though the data set is tiny) that fB declines with stellar age (Linsky and Saar 1987; Saar 1989). The magnetic field strength versus age diagram shows only scatter, implying that a reduction in the surface filling factor of active regions is the primary reason for the decline in stellar activity with time (at least for ages $\geq$ 0.3 Gyr). The decline of fB with time increases with time, but the functional form cannot be determined precisely with the available data. This result, too, is broadly consistent with theoretical models of angular momentum loss (e.g., Weber and Davis 1967; Kawaler 1988; Stepień 1988b).

Unlike the field strength, f exhibits a large range independent of color. Flare stars (listed by variable star name in Table 1) appear to have the highest filling factors. The f values of dM stars suggests that the large H$\alpha$ filling factors derived by Giampapa (1985) refer to the chromospheric level, where f is indeed much larger than in the photosphere (Saar *et al.* 1987). Filling factors estimated by Montesinos *et al.* (1987) are generally much smaller than those observed.

Due to the intimate connection between chromospheric, transition–region, and coronal heating and magnetic fields on the Sun, it is natural to search for similar relations in the

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stellar data. Marcy (1984), once again, lead the way, noting correlations between magnetic parameters, Ca II, and X-ray emission. Saar and Schrijver (1987) found $F_x \propto (fB)^{0.9}$ and $\Delta F_{CaII} \propto (fB)^{0.6}$ for fB $\leq$ 300 G (where $\Delta F_{CaII}$ is the residual Ca II flux; Schrijver 1983). Above fB = 300 G, they found $\Delta F_{CaII}$ was saturated. Schrijver *et al.* (1989) note that these correlations are consistent with relations derived for the Sun, and with flux-flux relations derived for stars. Saar (1988c) derived power law relations (with ranging from 0.6 and 0.3) between fB and ultraviolet C IV, C II and O I line fluxes. Quillen *et al.* (1987) found that using X-ray luminosities and the simple coronal loop model of Golub (1983), they could roughly predict stellar magnetic fluxes to within a scaling factor. Estimates of the individual f and B values were rather poor, however. Stepień (1988a) and Jordan *et al.* (1987) have predicted expected magnetic flux – activity relations with some success (e.g., Jordan *et al.* estimate $F_X \propto (fB)^{1.8}$).

Groups have also searched for variability in stellar magnetic fields. Already in the first detections, Robinson *et al.* (1980) noted an apparent change on $\xi$ Boo A. Subsequent observations of the star by Marcy (1981) and Gondoin *et al.* (1986) yielded no evidence for magnetic line broadening at all, suggesting the star had a large range of magnetic variation (consistent with its Ca II emission). Rotational modulation of chromospheric and transition-region line fluxes with magnetic flux for the active dwarf $\xi$ Boo A support this picture (Saar *et al.* 1988). The same authors used simultaneous measurements of broadband linear polarization (which measures the net tangential component of the magnetic field) to permit a rough determination of the spatial distribution of active areas on the star. Some of this variability may have been due to changes in turbulence due to a "starpatch", as postulated by Toner and Gray (1988). On the other hand, it is possible (Bruning and Saar 1989) that the line bisector variations seen in both data sets are primarily a magnetic phenomenon (masked in the Toner and Gray analysis by the small $g_{eff}$ leverage: $g_{eff}$ = 1.0 versus $g_{eff} = 1.7$). More work on the interaction of magnetic fields and convection, and their role in line shapes, is certainly needed.

Other stars studied in detail so far seem not to vary as much as $\xi$ Boo A. Basri and Marcy (1988) found little change in their $\epsilon$ Eri spectra, confirming earlier evidence (Saar *et al.* 1986b) that its variability range is small. The active subgiant HD 17433 (see Bopp *et al.* 1989) has also been the subject of a campaign of simultaneous magnetic field, ultraviolet, and optical measurements (Ambruster and Saar 1988). Little variability was observed here either, except during a brief flare, when C IV emission increased significantly and there was weak evidence for a change in the magnetic parameters. Analysis of a similar campaign on the nearly pole–on flare star BD +26 730 is underway (Saar *et al.* 1989, in preparation).

## 4. A Comparison of $\epsilon$ Eri Measurements

As an example of progress in magnetic measurements and difficulties yet to be resolved, Table 2 shows a compilation of all magnetic measurements to date for the active K2 dwarf, $\epsilon$ Eri (HD 22049). All filling factors were adjusted (where possible) to the scale of Marcy (1984; $\overline{\gamma} = 34°$).

Clearly, there has been some evolution of opinion as to the magnetic activity on the star, and the situation is perhaps still not clearly resolved. Early approaches which did not include radiative transfer effects in the Zeeman process (Marcy 1984; Gray 1984) generally

Table 2: Comparison of Magnetic Measurements for the K2 dwarf, $\epsilon$ Eridani

| Reference | B (kG) | f (%) | fB (kG) | $\sqrt{\text{fB}}$ (kG) | $\lambda$ (nm) | # of lines g<1.3/g≥1.3 |
|---|---|---|---|---|---|---|
| Marcy (1984) | 1.17 | 67 | 0.78 | 0.96 | 617 | 1/1 |
| Gray (1984) | 1.9 | 36 | 0.69 | 1.14 | 640 | 10/6 |
| Saar *et al.* (1986b) | 1.9 | 13 | 0.25 | 0.69 | 617 | 3/2 |
| Saar (1988) | (if 2.0 | ≤5) | ≤0.10 | ≤0.40 | 2210 | 1/1 |
| Saar (1988) | 3.0: | 8: | 0.24: | 0.85: | 617 | 3/2 |
| Mathys and Solanki (1988) | 1.79–2.53 | 10–20 | (0.30) | 0.80 | 580–680 | (45 total) |
| Basri and Marcy (1988) | 1.0 | 35 | 0.35 | 0.59 | 846 | 1/1 |
| Mathys and Solanki (1989) | 2.09–2.96 | 10–20 | (0.35) | 0.94 | 580–680 | 33/9 |
| Marcy and Basri (1989) | 1.0 | 30 | 0.30 | 0.55 | 846 | 1/1 |
| Marcy and Basri (1989) | (if 3.0 | ≤10) | ≤0.30 | ≤0.95 | 2210 | 1/0 |
| | (if 1.0 | ≤30) | ≤0.30 | ≤0.55 | 2210 | 1/0 |

produced higher filling factors. As noted in section 2, this is likely due to the need for higher f to mimic the more "box–shaped" (due to increased line opacity) doppler core of the line, in the absence of a radiative transfer treatment (e.g., Saar 1988a). Introduction of a simple (Unno 1956) radiative transfer model led to smaller filling factors (Saar *et al.* 1986b; average of 11 measurements). Later, analysis of simultaneous optical and infrared spectra using the same model indicated lower fluxes (Saar 1988a). Indeed the optical results were uncertain enough (due to the low f value) to be deemed a non-detection. The completely different, regression analysis approach of Mathys and Solanki (1988) yields f and B values suprisingly similar to Saar *et al.* (1986b), though the later recalibration (Mathys and Solanki 1989) produced magnetic field strengths larger than indicated by infrared data (Saar 1988a; Marcy and Basri 1989). The Marcy and Basri (1989) analysis includes improved radiative transfer and explicit disk–integration, and results in lower B and larger f values, though fB is approximately the same. Thus, happily, there is now some convergence in the results for fB and $\sqrt{\text{f}}$B, but further work is needed to understand the why the separate f and B values differ. As noted in section 3, part of the disagreement in the derived f and B for $\epsilon$ Eri (and other targets) is probably due to the differing wavelengths and lines used, and part is due to the different analysis techniques themseves. Finally, one should remember that even in the best of cases, it is difficult to uniquely separate f and B (Gray 1984).

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## 5. Some Future Directions

What questions need to be addressed in stellar magnetic research over the next few years? First, many more magnetic measurements are needed. Determinations for more M dwarfs and RS CVn's as well as first detections of T Tauri's and F dwarfs would be especially useful. Far infrared measurements of the 12 micron lines (Deming *et al.* 1988) and near infrared measurements using cryogenic echelles now being built will substantially increase the accuracy of the f and B determinations. More multiwavelength campaigns would help probe the temporal dependence of magnetic flux and its relationship to atmospheric heating. Observations of broadband linear (Huovelin *et al.* 1988) and circular (Kemp *et al.* 1987) polarization should help unravel some information on the spatial structure of the fields, given suitable theory (Landi Degl'Innocenti 1982; Mürset *et al.* 1988). It may be possible to obtain further spatial information through Stokes V (Donati *et al.* 1989) or Stokes I Doppler imaging (the spot profile "bumps" will be broader in high $g_{eff}$ lines). Multiline observations and analyses (e.g., Mathys and Solanki 1989) should be further explored and refined. Observations over stellar cycle timescales should prove enlightening.

Several important assumptions have been made in stellar Zeeman analyses to date, and it is important to reassess them. The assumption of identical magnetic and non–magnetic atmospheres should eventually be discarded in favor of realistic flux–tube atmospheres for the magnetic component. Infrared data may require a third, umbral component (Sun *et al.* 1987). Turbulent and convective properties in magnetic regions appear to differ from the quiet solar atmosphere (Livingston 1982) and may in stars as well (Toner and Gray 1988). This should be investigated in detail. Magneto–optical effects should also be included in the modeling (Landolfi *et al.* 1989).

In summary, much has been learned in the past ten years about magnetic fields on solar–like stars, and their relationship to both basic stellar properties and to magnetically–related activity. Only the surface of the subject has been really explored, however, and *many* questions remain. Hopefully, ten years hence, many of these will also have been answered, bringing better understanding of the stellar "activity" phenomenon.

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