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Searching for links between magnetic fields and stellar evolution[*]

II. The evolution of magnetic fields as revealed by observations of Ap stars in open clusters and associations

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ABSTRACT

Context. The evolution of magnetic fields in Ap stars during the main sequence phase is presently mostly unconstrained by observation because of the difficulty of assigning accurate ages to known field Ap stars.

Aims. We are carrying out a large survey of magnetic fields in cluster Ap stars with the goal of obtaining a sample of these stars with well-determined ages. In this paper we analyse the information available from the survey as it currently stands.

Methods. We select from the available observational sample the stars that are probably (1) cluster or association members and (2) magnetic Ap stars. For the stars in this subsample we determine the fundamental parameters $T_{\rm eff}$, $L/L_\odot$, and $M/M_\odot$. With these data and the cluster ages we assign both absolute age and fractional age (the fraction of the main sequence lifetime completed). For this purpose we have derived new bolometric corrections for Ap stars.

Results. Magnetic fields are present at the surfaces of Ap stars from the ZAMS to the TAMS. Statistically for the stars with $M > 3\,M_\odot$ the fields decline with advancing age approximately as expected from flux conservation together with increased stellar radius, or perhaps even faster than this rate, on a time scale of about 3×10^7 yr. In contrast, lower mass stars show no compelling evidence for field decrease even on a timescale of several times 10^8 yr.

Conclusions. Study of magnetic cluster stars is now a powerful tool for obtaining constraints on evolution of Ap stars through the main sequence. Enlarging the sample of known cluster magnetic stars, and obtaining more precise rms fields, will help to clarify the results obtained so far. Further field observations are in progress.

Key words. stars: magnetic fields – stars: chemically peculiar – open clusters and associations: general – stars: fundamental parameters – stars: evolution – stars: early-type

1. Introduction

A small fraction, of the order of 10%, of the main sequence A and B stars show obvious signs of anomalous atmospheric chemical abundances in their spectra. The distinctive surface abundances of these stars have never found a satisfactory explanation as the result of formation in a chemically anomalous region of the galaxy, or as the product of internal nuclear evolution or of surface nuclear reactions. Instead, it is now generally thought that these stars have undergone chemical fractionation as a result of competition between gravitationally driven diffusion of trace elements, radiative levitation of some of these elements, and interaction of the separation process with mass loss and with mixing processes such as convection and turbulence (Michaud 1970; Babel 1992; Richer et al. 2000).

One of the major groups of these main sequence stars with anomalous atmospheric abundances is the "magnetic peculiar A" stars, which are commonly referred to as Ap stars. These stars all have large-scale (global) magnetic fields with strengths in the range of a few hundred to about 30 000 G. Their distinctive chemical anomalies appear to vary fairly systematically with $T_{\rm eff}$; the cooler ones (Ap SrCrEu stars with $7000 \leq T_{\rm eff} \leq 10\,000$ K) show strong overabundances of Sr, Cr, and rare earths, while hotter ones (Ap Si and Ap He-wk stars, with $10\,000 \leq T_{\rm eff} \leq 15\,000$) typically show strong Si overabundance and He underabundance. The hottest Ap stars (the He-str stars with $T_{\rm eff} \sim 20\,000$ K) have overabundant He. The broad features of which elements are expected to be over or underabundant as a

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[*] Tables 2 and 3 are only available in electronic form at http://www.aanda.org

function of T_{eff} are understood – to some extent – as a result of the competing processes mentioned above.

The magnetic Ap stars are also usually periodically variable, in light, spectrum, and magnetic field. The periods range from about half a day to decades without any obvious difference in the phenomena of variation, and the periods observed are inversely correlated with $v_{\text{e}} \sin i$. It is clear from these facts that the period of variation must be the rotation period. This variability implies that the field and the chemical abundances are non-uniformly distributed over the surface of an Ap star. The rotation axis is not an axis of symmetry, so that the observer sees a configuration that changes as the star rotates.

Much is now understood about these magnetic Ap stars. Simple models of field structure indicate that almost all magnetic Ap stars have roughly dipolar field structures. The axis of (approximate) symmetry of the field is usually oblique to the rotation axis, so that the line-of-sight component of the field varies roughly sinusoidally with time (e.g. Mestel & Landstreet 2005). Often several chemical elements are distributed quite non-uniformly over the surface, and low-resolution models (e.g. Strasser et al. 2001) of the distributions are available for a number of Ap stars. It is now becoming possible to deduce quite detailed maps of both field structure and element distribution for a few stars for which the variations have been observed spectroscopically in all four Stokes parameters (Kochukhov et al. 2004; Kochukhov 2004).

The lack of any energetically significant convection zone near the surface of all but the coolest Ap stars, and the absence of a direct correlation between rotation rate and field strength (as is found in solar-type stars, see Mestel & Landstreet 2005) has led to general agreement that the magnetic fields observed are probably fossil fields retained from an earlier evolution stage, a situation which is possible because of the extremely long decay times.

Numerical calculations have suggested how a fossil magnetic field might evolve with time during the long main sequence phase as a consequence of ohmic decay, changes in stellar structure, and internal circulation currents (Moss 2001), but the relationship of such computations to observed stars is still very unclear. Braithwaite & Spruit (2004) and Braithwaite & Nordlund (2006) have shown that the combination of the Cowling instability with rotation and meridional circulation may lead to major reorganisation of the internal flux. Similarly, theoretical computations suggest how atmospheric chemical peculiarities might develop and evolve with time through the main sequence phase, but so far these computations have not succeeded in reproducing observed abundance patterns in detail (Babel & Michaud 1991). Observational data on fields and abundances in a sample of Ap stars for which the absolute age and fraction of the main sequence evolution completed (we will call this the "fractional age") is known would be extremely valuable to test theoretical ideas and provide clues to important processes.

At present, we have few constraints from observations about the evolution of either fields or atmospheric chemistry with time before or through the main sequence phase. Ap stars have somewhat anomalous, and variable, energy distributions, which make it difficult to place these stars in the HR diagram precisely enough to obtain accurate ages for individual stars from comparison with theoretical evolution tracks. The problem is compounded by the fact that the atmospheric chemistry does not readily reveal the bulk stellar abundances, and so we do not even know how to choose the metallicity used in computing the evolution tracks which are to be compared with observations. As a result, until recently very few Ap star ages could be securely

determined. Age determinations of field stars are barely accurate enough to determine whether a given star is in the first half or second half of its main sequence life. (This problem is discussed in more detail by Bagnulo et al. 2006, hereafter Paper I).

The problem of obtaining an accurate age for an Ap star is considerably simpler if the star is a member of a cluster. Since the age of a cluster can be determined typically to somewhat better than ±0.2 dex (by isochrone fitting to the main sequence turnoff or to low-mass stars evolving to the zero-age main sequence, or by studies of Li depletion in low-mass stars), we can use the cluster age to determine the fraction of the main sequence lifetime which has elapsed. For a star in a cluster whose age is small compared to the star's main sequence lifetime, this evolutionary age is relatively precise. On the other hand, fractional age resolution is not very good for the second half of a star's main sequence lifetime (see Paper I, and Sect. 5.1).

It has recently become practical to obtain magnetic measurements of a statistically interesting number of stars in clusters. In Paper I we have reported the first large scale observational survey of magnetic fields in peculiar (and normal) A and B stars in open clusters and associations. This survey has been motivated by our desire to obtain empirical information about the evolution of magnetic fields and atmospheric chemistry in middle main sequence stars as these stars evolve from the zero-age main sequence (ZAMS) to the terminal-age main sequence (TAMS).

In the present paper, we draw some first conclusions from this (ongoing) survey about the evolution of magnetic fields from ZAMS to TAMS. In Sect. 2 we will examine and select all the stars that are probable cluster or association members and probable Ap stars, and for which magnetic measurements are available (either from our own survey or previous works). In Sect. 3 we determine the fundamental parameters T_{eff} and $L/L_{\odot}$ for these stars. Combining this information with suitable evolution tracks, we derive stellar masses and fractional ages for the stars of the sample in Sect. 4. In Sect. 5 we comment on the characteristics of our sample (distributions of age, mass, and magnetic field). Finally, in Sect. 6, we consider what these data reveal about the evolution of surface fields from the ZAMS to the TAMS. Section 7 summarises our conclusions.

2. Cluster membership of stars with magnetic data

In Paper I we presented new magnetic measurements of many candidate or confirmed Ap stars that may be members of clusters or associations (as well as a large number of measurements of probable cluster members that are not Ap stars). From these data and from field measurements in the literature, we have assembled a database of all probable Ap cluster/association members for which usefully precise magnetic field measurements are available. This database is presented in Table 2; the content will be explained below. In this section we analyze more critically the evidence for cluster membership, and for magnetic Ap nature, for the stars in Table 2. This will enable us to remove any stars for which either membership or Ap nature is doubtful from the stellar sample we finally analyse.

2.1. Assessing cluster membership

Cluster (or association) membership may be tested in several ways. Traditionally, membership is tested by plotting a colour-magnitude diagram for the cluster. Stars which lie too far above or below the cluster main sequence are rejected as members. Previous studies (e.g. North 1993; Maitzen et al. 1988) have

shown that Ap stars that are probable cluster members lie reasonably close (in observational HR diagrams, for example Johnson V magnitude as a function of colour $B - V$) to the main sequence of normal stars, although Ap stars sometimes deviate from the normal main sequence by somewhat more than normal cluster members. In general, the stars in Table 2 are sufficiently close to the main sequences of the clusters to which they may belong that this criterion does not immediately eliminate any of the stars in the table from membership. The HR diagram test of membership will be applied again later in this paper when the provisionally accepted cluster members are placed in the theoretical ($L/L_\odot$ vs. $T_{\rm eff}$) HR diagrams of their specific clusters.

Comparisons of measured parallaxes and proper motions of candidate stars with the average values for the cluster provide powerful tests of membership. The situation for data on both these properties has improved dramatically as a result of the Hipparcos mission (ESA 1997). This mission led to measurements of parallaxes and proper motions of about 120 000 stars, with typical errors of the order of 1 mas and 1 mas yr^{-1}, so that accurate ($\pm 25\%$) parallaxes are available for tens of thousands of stars out to a distance of about 300 pc. Observations of a much larger number of stars than that included in the Hipparcos parallax programme, using the star mapper instrument on the satellite, led to the Tycho (or Tycho-1) Catalogue (Høg et al. 1997) of proper motions of about 1 million stars, with accuracies of the order of 10 mas yr^{-1} (at $V = 9$) in both coordinates. The Tycho data were combined with positions from the Astrographic Catalogue to produce the Tycho Reference Catalogue (Høg et al. 1998), with proper motions of 990 000 stars accurate to about 2.5 mas yr^{-1}, and an extension of this project led to the Tycho-2 catalogue (Høg et al. 2000a,b), which combines data from the Hipparcos mission with those from many ground-based position catalogues to provide proper motions for about 2.5 million stars with typical proper motion standard errors of 2.5 mas yr^{-1}.

These data have been used by several groups for important studies of membership in clusters and OB associations. Robichon et al. (1999) determined mean distances, mean proper motions, and membership of Hipparcos parallax stars located near the centres of 18 clusters less than 500 pc away. Membership was derived from assessment of position, parallax, proper motion, and photometry. The work led to list of accepted members of each cluster; the absence of a Hipparcos parallax star from one of these lists presumably means that it was not accepted by Robichon et al. (1999) as a member of the corresponding cluster.

Baumgardt et al. (2000) derived mean distances and motions for 205 open clusters using Hipparcos parallax stars found in these clusters. Membership of Hipparcos parallax stars in these clusters was assessed using photometry, radial velocity, position, and proper motion data from the ground and from Hipparcos, together with some additional proper motions from the Tycho Reference Catalogue. Baumgardt et al. report both an overall membership assessment (member, possible member, non-member) and a proper motion membership probability.

The Hipparcos data were used by de Zeeuw et al. (1999) to assess membership of Hipparcos parallax stars in nearby OB associations. Several of the associations included in their study (Sco OB2, α Per, Collinder 121, Trumpler 10, etc.) are represented in our survey.

The release of the Tycho-2 proper motion catalogue made possible further large-scale studies of membership in open clusters. Dias et al. (2001, 2002) have determined mean proper motions of more than 200 clusters, many of which were not included in the studies discussed above, using in most cases several tens of stars per cluster. They have derived membership probabilities based only on the proper motions, which are tightly clustered for an open cluster but (usually) have a considerably larger dispersion for field stars.

Kharchenko et al. (2005) created a database of possible stellar members of some 520 clusters. They use a combination of Tycho-2, Hipparcos, and ground-based data to determine average cluster parameters (distance, reddening, proper motions, etc.). Memberships of individual stars in their database are assessed using all available information (position, photometry, proper motions, etc.), and then cluster ages are derived using the positions of the brightest stars relative to theoretical isochrones.

We have used the studies discussed above together with the actual astrometric data to assess the probability that the stars of Table 2 are actually cluster members. Wherever possible we have adopted the consensus of the previous membership studies. However, the individual determinations of membership probability differ somewhat from one study to another, both because a study may be based either on Hipparcos or Tycho-2 proper motions, and because the stars selected as definite cluster members, from which the cluster mean motions are deduced, varies from one study to another. Note in particular that Kharchenko et al. (2005) often assign membership probabilities based on proper motions that are significantly smaller than those found in any of the other studies (e.g. for HD 21699 in the α Per cluster, de Zeeuw et al. 1999 give $P = 0.84$ and Robichon et al. 1999 consider the star to be a definite member, but Kharchenko et al. 2005, give $P_\mu = 0.25$ as the proper motion membership probability). We generally consider that the Kharchenko et al. data support membership if their proper motion membership probability is larger than about 0.15 (their 2-σ limit).

In a few cases the literature conclusions on membership are discordant, or no membership probability based on recent astrometry is available. In these cases we have compared proper motions and parallaxes, when available, to the cluster means, and assessed membership ourselves. We consider a star to be a probable member if both proper motions (and parallax if available) are less than 2.5σ from the mean cluster values, and a definite member if all are within 1σ of the cluster means. If no accurate astrometry is available, but the star is within the bounds of the cluster and has colour and magnitude consistent with membership, we have generally assumed that the star is a probable member.

In the particular case of stars in Ori OB1a, OB1b, and OB1c, the members of the association are not well separated in proper motion from the field. We have adopted the criterion of Brown et al. (1999) for proper motion membership, and have tested apparent brightness and parallax for consistency with membership as far as possible. However, membership of stars in these regions of Ori OB1 is uncertain (see discussion in de Zeeuw et al. 1999). In Ori OB1d, where background obscuration virtually eliminates background stars, we can be more confident of membership.

The results of this exercise are reported in Table 2 in the four columns under the general heading "cluster membership". In these columns we report our conclusions as to membership probability considering only parallax ("π"), proper motions ("μ"), location in the HR diagram relative to the cluster isochrone ("phot", see below), and overall ("member?"), using a four-level scale of "y" = member, "p" = probable member, "?" = probably non-member, and "n" = non-member. The "member?" column also contains references to earlier membership studies. Notes at the bottom of Table 2, signaled by "*" in the "member?" column, discuss particular problems with deciding on cluster or association membership.

2.2. Assessing spectral type and chemical peculiarity

We next turn to methods to determine whether a particular star is an Ap star of the type in which a magnetic field is almost always detected when sufficiently precise field measurements are obtained (Aurière et al. 2004). The reason that this is important is that the field measurement errors achieved in general in Paper I are only small enough to detect fields in about half of the magnetic Ap stars we have observed, and since we will want to examine the distribution of field strengths in magnetic Ap stars as a function of age and mass, it is important to consider both the Ap stars in which we detect a field and those in which no field has yet been found.

Several criteria are available that can indicate Ap type spectra. Spectral classification can provide an important test of Ap characteristics. For many of the stars in our sample, one or more classification spectra are available. Classification surveys of a number of clusters with the goal of identifying Ap stars have been reported by Hartoog (1976, 1977) and Abt (1979) and references therein. Spectroscopic studies of single clusters have also yielded many spectroscopic classifications (these may be found in the WEBDA database of open cluster data at `http://www.univie.ac.at/webda/`, described by Mermilliod & Paunzen 2003; on the SIMBAD database of stellar data at `http://cdsweb.u-strasbg.fr/`; and in the compilation of data on Ap stars in clusters by Renson 1992). Unfortunately, since the classification criteria for Ap spectra are not always unambiguous, and may not be clearly visible in low dispersion, low signal-to-noise classifications spectra, the available classifications of a star are frequently contradictory as to whether the star is actually an Ap.

Another valuable criterion for Ap type is furnished by the Δa photometry system developed by Maitzen and collaborators (e.g., Maitzen 1993), and by the Z index which can be computed for stars for which Geneva photometry is available (e.g. Cramer & Maeder 1979, 1980; Hauck & North 1993). Both of these photometric indices are sensitive to the broad energy depression at about 5200 Å which has been found to be a robust indicator of Ap spectral type in roughly the temperature range 8000−14 000 K (e.g., Kupka et al. 2003). Both these photometry systems provide a large amount of information about cluster Ap stars; Maitzen's group has surveyed a large number of clusters for Ap stars, and Geneva data exist for many cluster stars as well (see `http://obswww.unige.ch/gcpd/`).

A third useful criterion of Ap type is the detection of periodic photometric variability of the type frequently found in Ap stars. The catalogues of Renson & Catalano (2001) provide very useful summaries of the many particular studies found in the literature. We have searched the Hipparcos catalogue (ESA 1997) for new photometric variables, although this provided only a very small number of variables that had not already been identified by ground-based photometry.

Finally, the definite detection of a magnetic field is a clear indication of Ap classification.

We have summarized our view as to whether the star is an Ap star in the four columns under the general heading "Ap star" of Table 2. We have used the same scale adopted to evaluate the membership ("y", "p", "?", and "n"; a blank space denotes absence of information). In the column labelled "Sp" we report the conclusions obtained from the available spectral classifications. We have adopted the criterion that Δa photometry definitely supports Ap classification if $\Delta a \geq 0.014$, and that Geneva photometry supports Ap classification if $Z \leq -0.016$ (within the temperature range where photometric detection of peculiarity is valid). The conclusions are given in column Δa, Z. The degree of support for Ap classification from photometric variability is summarized in the column labelled "var". The column "mag fld" reports our assessment as to whether the star has a detected field, and also provides references to field measurements. From this column it will be clear that a large fraction of all field measurements of open cluster stars have been carried out in the course of our survey (Paper I; measurements denoted by "F" in the "mag fld" column); the only substantial previous surveys have been restricted to the Sco OB2 and Ori OB1 associations. Finally, an overall assessment as to whether the balance of evidence supports an Ap classification is provided in the last column "Ap?" of Table 2. Note that a significant number of stars are given a "y" or "p" even though no magnetic field has yet been detected in them.

3. Determining the physical properties of cluster Ap stars

With the analysis described in the previous section we have identified a subsample of about 90 stars that are definite or probable ("y" or "p") cluster/association members, definite or probable Ap stars, and for which at least one magnetic field measurement is available (regardless of whether a field has been detected or not).

For each star of this subsample we need to determine age and mass. These will be calculated in Sect. 4, using suitable evolutionary models, assuming that we know the cluster age, the star's temperature and the star's luminosity. In this section we describe how we determine these physical properties. Furthermore, to each star we associate a value representative of the star's field strength.

3.1. Cluster ages

Because our essential goal is to provide a sample of magnetic measurements of Ap stars of reasonably well known ages, the ages of the clusters to which the stars in our sample belong must be critically discussed. Cluster ages for a large sample of clusters, including essentially all those considered here, are available from the WEBDA database (Mermilliod & Paunzen 2003). Ages in this database are derived from a survey of literature values carried out by Loktin et al. (2001). These ages are in most cases based on isochrone fits to dereddened cluster main sequences.

Another extensive study of ages has been carried out by Kharchenko et al. (2005) based on databases of potential cluster members generated from large observational resources such as the Hipparcos and Tycho databases. The ages of Kharchenko et al are derived by fitting isochrones individually to the most massive main sequence stars found among their own selection of cluster members, rather than by isochrone fitting to the upper end of the cluster main sequence as a whole. The ages of Kharchenko et al. are based in many cases on only one star per cluster. Nevertheless, these ages are in good overall agreement with those of Loktin et al. (2001): Kharchenko et al. have shown that (excluding a small number of clusters where significant differences in choices of cluster members lead to substantially different ages) the ages from these two sources exhibit an rms difference in cluster age ($\log t$) of about $\sigma = 0.20$ dex. Note that neither WEBDA nor Kharchenko et al. provide uncertainties for the assigned cluster ages; one of our main tasks is to estimate these uncertainties.

We have included these ages in our own database of possible cluster Ap stars, together with many recent determinations from the literature. After examining the available recent age determinations, we have selected what seems to us to be an appropriate age for each cluster, and assigned an uncertainty to this value. In assigning ages, we have considered both accuracy claimed by individual age studies and the overall agreement of various studies. For a number of the clusters of interest, several age determinations are available, and in many cases these are reasonably concordant. For such clusters we have assigned an uncertainty which in the best cases is taken to be ±0.1 dex. For clusters with fewer age determinations, or more discordant ones, we have generally assigned an age and an uncertainty which approximately cover the range of the best age determinations within the range 0.1−0.2 dex. For clusters for which no ages seem to be available, we have mostly used the ages from WEBDA (those of Loktin et al. 2001), and assumed an uncertainty of ±0.2 dex, in agreement with the overall concordance between their results and those of Kharchenko et al. (2005).

In addition, several of the probable Ap stars we have observed turn out to be near the TAMS in the HR diagram. By requiring the cluster isochrone to intersect the error oval of such a relatively evolved star, in some cases we are able to constrain the cluster age more precisely than by using ages obtained from the literature; we obtain uncertainties that may be as small as 0.05 dex. In these cases we have adopted the cluster ages and uncertainties from our isochrone fits. The procedure adopted is discussed further in Sect. 4.

3.2. Stellar effective temperatures

3.2.1. Photometry

For stars for which Geneva photometry is available, we have used the FORTRAN code described by Künzli et al. (1997), which is available from CDS, to determine a first effective temperature based on the calibration for normal (metallicity 0) stars. This procedure requires supplying a colour excess $E(B_2 - V_1)$. We have obtained this quantity from the Johnson colour excess $E(B - V)$ (either for the cluster as a whole, as reported by WEBDA or a particular study, or for the star individually after de-reddening in the Johnson $U - B, B - V$ colour−colour diagram), using the relationship $E(B_2 - V_1) \approx 0.88E(B - V)$ (Hauck & North 1993). When the colours of a star are near the boundary between different methods of temperature determination (e.g. XY or $pTpG$) we have computed the temperature using both methods. The effective temperature determined using normal star calibrations is then corrected to the Ap stars temperature scale as described by Hauck & Künzli (1996). For stars classified as He-weak or He-strong, we have followed the suggestion of Hauck & North to adopt the temperature produced by the Künzli et al. FORTRAN programme without correction.

For stars for which Strömgren $uvby\beta$ photometry is available, we have computed an effective temperature using the FORTRAN code "UVBYBETANEW" of Napiwotzki, Schönberner & Wenske (1993), which is in turn based mainly on the calibrations of Moon & Dworetsky (1985). When the star has colours near the boundary of different calibrations (for example near A0), we have computed the temperature using both calibrations. These raw temperatures have been corrected for Ap stars as described by Stępień & Dominiczak (1989). Again, we have not corrected temperatures of He-weak or He-strong stars from the values produced by the programme.

If only Johnson UBV photometry is available, we have de-reddened the star to the main sequence line in the $U - B, B - V$ colour−colour diagram, and then corrected the value following the same prescription Stępień & Dominiczak (1989) use for Strömgren photometry.

3.2.2. Adopted temperatures and uncertainties

For most of the stars in our sample, both $uvby\beta$ and Geneva photometry are available, and we have adopted a suitable average of the computed temperatures. We find that temperature determination using the two different kinds of photometry, and sometimes more than one method of temperature determination, have a scatter of the order of 2−300 K. Since both methods have been calibrated using mostly the same "fundamental" stars, this is certainly a lower bound to the actual uncertainty of the results.

The global temperature calibration for Ap stars is still not very satisfactory. *No* Ap stars have a fundamental calibration (based on integrated flux and angular diameter measurements); all calibrations are based on corrections to the normal star calibrations using a small sample of stars for which effective temperatures have been determined using the infrared flux method (e.g. Lanz 1985; Megessier 1988; Monier 1992), fitting energy distributions (e.g. Adelman et al. 1995), or the reasonable but uncertain correction method used by Stępień & Dominiczak (1989), who estimate the uncertainty of their individual values of T_{eff} to be 6−700 K. Furthermore, the accuracy of these corrections has recently been called into question by the results of Khan & Shulyak (2006), who have calculated model atmospheres with enhanced metal content (so far generally scaled solar abundances) *and* the effects of a magnetic field. They find that, even with a rather large field and 10x solar abundances, the Paschen continuum slope is very similar to that of a solar composition, non-magnetic star, and that the calibrations for *normal* stars recover approximately the assumed effective temperature.

Finally, there are certainly substantial star-to-star variations in atmospheric composition, which make any calibration suitable only on average, and not exact for any particular star. This is clearly seen in the scatter shown in Figs. 9 and 10 of Napiwotzki et al. (1993), where individual calibration points are shown along with the mean calibration curves adopted.

Considering the internal uncertainties in determining the effective temperature of a magnetic Ap star, and the further global uncertainties, we adopt a uniform uncertainty of ±0.02 dex (roughly ±500 K) for our temperatures, and even this value may be somewhat optimistic.

3.3. Stellar luminosities

3.3.1. Cluster distances

In the case of the nearest clusters (out to about 300 pc) accurate distances are best derived using Hipparcos parallaxes averaged over known members. Distances obtained in this way generally have uncertainties of about 0.2 mag or less in both absolute and apparent distance modulus (cf. Robichon et al. 1999).

For more distant clusters, the most accurate distances are obtained by isochrone fitting to dereddened cluster main sequences. For a number of nearby clusters, it is possible to compare distances obtained from isochrone fitting with those obtained from parallaxes (Robichon et al. 1999). This comparison, and comparison of different determinations using isochrone fitting, allow us to estimate that a typical uncertainty for the more distant clusters is roughly 0.2 mag in distance modulus, similar

to the value for the nearer clusters. In almost all cases the uncertainty in the distance modulus is dominated by the uncertainty in the distance rather than that in the reddening.

We have adopted the cluster or association distances for all of the stars in our sample, rather than individual distances even when these are available from parallax observations; on average the mean cluster distances from parallaxes have standard errors that are smaller by a factor of two or more than those of individual cluster stars. Of course, for stars in more distant clusters, the best available distance is the photometric distance modulus of the cluster as a whole.

3.3.2. Bolometric corrections

The results of Lanz (1984) strongly suggest that the bolometric corrections for magnetic Ap stars are different from those of normal stars. Since we have effective temperatures obtained from several photometric systems, we need bolometric corrections suitable for Ap stars as a function of effective temperature, rather than as a function of a specific photometric index (as usually provided). We have re-examined the available data to obtain accurate bolometric corrections in the form we need.

For this purpose, we have collected all the magnetic Ap stars for which integrated fluxes are available. Bolometric corrections for magnetic Ap stars have been reported by Lanz (1984), while Stępień & Dominiczak (1989), and Monier (1992) have provided further integrated fluxes. Almost all the stars discussed by Lanz are rather hot (Si and He-pec stars), and most have distances between 100 and 250 pc, so the observed fluxes are significantly affected by interstellar absorption. Except for three cool Ap stars in his sample, the fluxes reported by Lanz have been corrected approximately for extinction. The integrated fluxes reported by Stępień & Dominiczak do not seem to be corrected for absorption, even though this sample includes some (hot) Ap stars as far away as 150−180 pc, for which interstellar absorption is probably significant. (However, for the five stars in common with Lanz's sample, with distances in the range 90−180 pc, the average difference in integrated flux between the two studies is only about 4%.) The three values of integrated flux reported by Monier are not corrected for interstellar absorption, but all refer to stars located between 50 and 100 pc from the Sun, and any corrections should be at most a few percent.

We have assigned effective temperatures to the stars for which integrated fluxes are available using the same methods discussed above, and computed the bolometric corrections for the stars of Stępień & Dominiczak (1989), and Monier (1992) using the first equation in Lanz (1984). Because of the lack of correction for interstellar extinction, we have discarded all the stars from Stępień and Dominiczak's sample that have $T_{\rm eff} > 11\,000$ K. However, we verified that including them did not significantly change the results.

Figure 1 shows the resulting data set of bolometric corrections as a function of $T_{\rm eff}$ together with our best fit (lower curve) to the Ap star mean correction $BC_{\rm Ap}(T_{\rm eff})$:

$$BC_{\rm Ap}(T_{\rm eff}) = -4.891 + 15.147\,\theta - 11.517\theta^2, \qquad (1)$$

where $\theta = 5040.0/T_{\rm eff}$.

We also plot the variation of bolometric correction $BC_{\rm n}(T_{\rm eff})$ with $T_{\rm eff}$ for normal stars, using (upper curve) Schmidt-Kaler's values (Lang 1992), and Balona's 1994 fit (middle curve) to the fundamental data of Code et al. (1976) and of Malagnini et al. (1986).
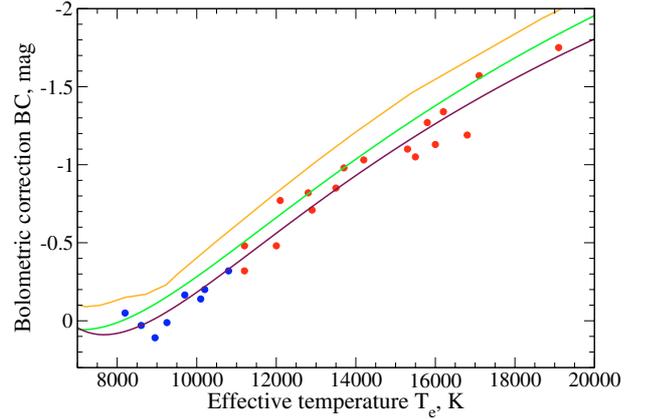

Fig. 1. Bolometric corrections using integrated fluxes of magnetic Ap stars from Lanz (1984, red circles), Stępień & Dominiczak (1989, blue circles), and Monier (1992, blue circles) as functions of effective temperature. The upper solid line is the BC for normal stars from Schmidt-Kaler (Lang 1992); the middle line is Balona's (1994) fit to the fundamental data of Code et al. (1976) and Malagnini et al. (1986), and the lowest curve is a polynomial fit to the Ap data as described in the text.

Equation (1) describes the available bolometric corrections for Ap stars between about 7500 and 18 000 K. The rms deviation of the individual Ap star data around this curve is about 0.1 mag, considerably larger than that for normal stars around Balona's fit. This reflects both larger uncertainty in both coordinates, and probably also real and substantial star-to-star variations in the bolometric correction for these rather diverse objects.

It is clear that the values of Code et al. (1976) and Malagnini et al. (1986) are the most accurate data for the normal star bolometric correction $BC_{\rm n}(T_{\rm eff})$. Note that the BC values of Schmidt-Kaler, which are sometimes used for Ap stars (e.g., Hubrig et al. 2000) are systematically too negative compared to the more accurate data of Code et al. (1976) and Malagnini et al. (1986), and differ from the calibration adopted here by about 0.25 mag.

With typical distance modulus uncertainties of 0.2 mag, an uncertainty in the bolometric correction of about 0.1 mag, and reddening uncertainties of less than 0.1 mag, we estimate that the typical uncertainty in $M_{\rm bol}$ is about 0.25 mag, corresponding to an uncertainty in $\log(L/L_\odot)$ of about 0.1 dex. We adopt this value as the standard error of luminosity values in Table 3.

The apparent V magnitudes and values of $T_{\rm eff}$ of stars in our subsample have been used, together with the bolometric corrections of Eq. (1), to determine the stellar luminosities $L/L_\odot$. For most clusters we used the mean cluster reddening as tabulated by WEBDA or one of the specific references. However, we found that using the mean cluster or association reddening led in several cases to stars appearing seriously underluminous compared to the position of the expected main sequence. In these cases we have determined the reddening $E(B-V)$ directly for each star from its position in a colour-colour diagram, and used these values (with $A_V = 3.1E(B-V)$) to determine the absolute magnitude of each individual star. This was necessary for Ori OB1, the Upper Sco region of Sco OB2, and for NGC 2244, 2516, and 6193. In general, this led to a much tighter observed main sequence.

3.4. Magnetic field

Magnetic fields of Ap stars may be determined through the measurement of the Zeeman splitting of spectral lines in a simple intensity spectrum, or through the analysis of circular polarization of spectral lines. The former method, exploited by Mathys et al. (1997), gives the average of the magnetic field modulus over the visible stellar disk. Zeeman splitting may be detected only under quite special circumstances, i.e., $v_e \sin i$ at most a few km s^{-1}, and field strength at least 2 kG, that are not met in most of the known magnetic Ap stars. Circular polarization measurements are generally a far more sensitive and broadly-applicable method of field detection, and give the so called *mean longitudinal magnetic field* $\langle B_z \rangle$, i.e., the component of the magnetic field along the line of sight averaged over the stellar disk.

Because of geometrical reasons, both mean field modulus and mean longitudinal field change as the star rotates. In particular, a $\langle B_z \rangle$ measurement could be consistent with zero even in a star that possesses a strong magnetic field. For the present work we need to associate to each star a field value that is representative of the stellar field strength.

For the few stars in our sample for which the full variation of the longitudinal field $\langle B_z \rangle$ is known and has been fitted with a sine wave of the form

$$B_z(t) = B_0 + B_1 \sin(\omega t - \phi_0), \tag{2}$$

we give the root-mean square longitudinal field strength $B_{rms} = (B_0^2 + B_1^2/2)^{1/2}$. For the great majority of the stars in the table, however, only one measurement, or at most a few scattered measurements, are available. In this case we simply report the raw value of B_{rms} averaged either over all available field measurements, or over the subset with the smallest uncertainties. Since most of the stars in this table have only one or two measurements, this is clearly not a very accurate statistic for an individual star, but averaged over our sample we believe that it already provides a significant amount of useful information.

When the individual field strength values used in the computation of B_{rms} are taken from Paper I, the value derived from the full spectrum is used if the measurement detects a field of less than 1 kG. However, as discussed in Paper I, the metal lines in FORS1 spectra yield an underestimate of the actual field strength for larger fields. Accordingly, for fields above 1 kG we have used only the field strength deduced from the Balmer lines.

Note that values of B_{rms} smaller than about 250 G typically should be read as indicating that the rms field estimated from the current data set is of the order of 200 G or less, regardless of the actual tabulated value.

4. Determining stellar masses and fractional ages

With effective temperatures and luminosities for almost all the stars in Table 3, we can now make comparison with evolution tracks to determine approximate masses and fractional ages (recall that the fractional age of a star is the fraction of its main sequence lifetime, measured from the ZAMS to the TAMS, that has already elapsed).

As discussed in Paper I, there is a spread of chemical compositions among nearby open clusters and associations of about a factor of 2.5 in metallicity, so that the accurate placement of a star on an evolution track requires either the use of evolution tracks appropriate to the particular star or stellar group, or consideration of the additional uncertainties introduced by variation of evolution tracks with metallicity. For studies of field stars, for which ages are determined from comparison of position in the HR diagram with evolution tracks, we showed in Paper I that very important age uncertainties arise from this effect, especially for stars in the first half of their main sequence lives. We also showed that it is only currently possible, with present uncertainties in T_{eff} and $L/L_\odot$, to discriminate between field stars in the first and second halves of their main sequence lifetimes. In contrast, uncertainty in bulk chemical composition does not introduce an important additional uncertainty into the determination of the stellar mass.

In this study we are determining stellar ages from the ages of the clusters to which the stars of our sample belong. The principal uncertainty for most of the available literature ages (0.1−0.2 dex in log t, as discussed above) arises from uncertainty as to where to place the cluster turnoff in the HR diagram, or in the appropriate models for pre-main sequence stars, rather than from uncertainty in the composition that should be adopted for the evolution tracks used for age determination. Consequently, we do not need to use composition-specific evolution tracks for mass and fractional age determination; we simply include the small uncertainty carried by this effect in our overall uncertainty estimates.

We then proceed as follows. Each star is placed in the HR diagram, using the values of T_{eff} and $L/L_\odot$ determined previously. Around each star we construct an approximate error oval, using the adopted uncertainties discussed above. The star positions are compared graphically, one cluster at a time, with evolution tracks for stellar masses in the range 1.7 to 9.0 $M/M_\odot$, and with the range of isochrones allowed by the adopted cluster or association age ±0.1 or 0.2 dex, using models computed by the Padova group (see Fagotto et al. 1994, and references therein) for $Z = 0.008$. 0.02 and 0.05. (These data are available in machine-readable form from CDS.) This comparison allows us to test the hypothesis that each star is actually a cluster member, and if it is, to determine its mass.

In almost all cases, the best overall agreement of individual star positions in the HR diagram with computed isochrones is found for $Z = 0.02$. This agreement persists even when the magnetic Ap stars are near the end of their main sequence evolution. This Z value is somewhat higher than the currently accepted best estimate for the Sun, which is about $Z = 0.012$ (Asplund et al. 2005). It is not clear whether our result indicates that $Z \approx 0.02$ is the most appropriate mean value for young stellar groups in the solar neighborhood, or whether this reflects a residual small systematic error in our transformation to the theoretical HR diagram (for example, modest underestimation of T_{eff} values as discussed above). Note that the significantly better fit of the $Z = 0.02$ models to our T_{eff}, $L/L_\odot$ data compared to the fit with $Z = 0.008$ models indicates that Padova models with $Z = 0.012$ would not fit our observational data as well as the $Z = 0.02$ models do. However, we cannot really test how well our data would fit isochrones computed with the new Asplund et al abundances, as these abundances have different ratios of light (CNO) elements to iron peak elements than are assumed in any available grid of evolutionary models.

We have tested the correctness of our computations, and assessed whether our assumed uncertainties are realistic, by looking at the results of such comparisons for nearby clusters and associations of well-determined membership, precisely known distance and moderate or small reddening. One such comparison is shown in Fig. 2 for the Sco OB2 association, using tracks for $Z = 0.02$. This figure confirms the general correctness of the transformations used in going from photometric data to the theoretical HR diagram, and shows that our assumed uncertainties
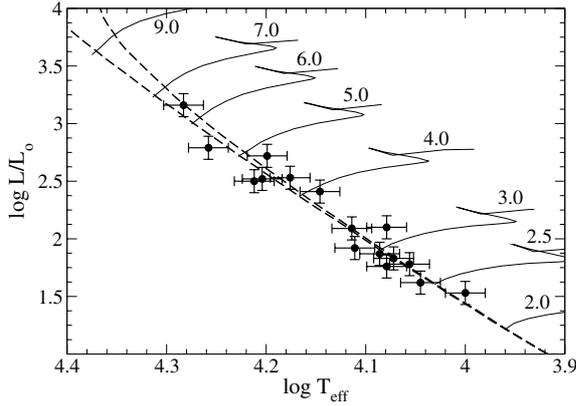
Fig. 2. Comparison of computed positions of stars in the Lower Centaurus-Lupus, Upper Centaurus-Crux, and Upper Sco regions of the Sco OB2 association with evolution tracks for stars of $Z = 0.02$ and $M = 2, 2.5, 3, 4, 5, 6, 7$, and $9\ M_\odot$, and with bracketting isochrones for log ages of 6.9 and 7.2, appropriate for these regions of the association.

give a reasonable description of the scatter about the $Z = 0.02$ isochrones.

In a few individual cases, typically in relatively distant or heavily reddened clusters or associations, we find significant disagreement between the deduced position of a star and the relevant isochrone, which strongly suggests that the star should be disqualified from cluster membership on photometric grounds. There are eight such cases in Ori OB1 and five in various clusters. All have been noted with "?" in the "phot" column of Table 2, and are omitted from Table 3. The other stars whose membership we have tested in this way, and whose positions are consistent with the isochrones to within better than 2σ are noted with "y" or "p" in this column. (We did not test stars whose cluster membership or Ap nature was rejected on other grounds.)

Placing the observed stars on the appropriate range of cluster isochrones allows us to provide a mass estimate for each. These masses have been derived using the $Z = 0.02$ evolution tracks. Uncertainties for these values, also reported in Table 3, arise primarily from the size of the error oval of each stellar position due to uncertainties in $T_{\rm eff}$ and $L/L_\odot$. A small contribution to the mass uncertainties is made by the uncertainty in the Z value appropriate for each cluster. For stars whose 1σ error ovals intersect the band of isochrones defined by the (inexact) age of the cluster, we find that the uncertainty in mass is typically about 5% of the actual mass value.

Our method of mass determination may be clarified using Fig. 2. Consider the star between the $M/M_\odot = 6$ and 7 evolution tracks (HD 125823). It may be seen directly from the figure that with this set of tracks and the indicated uncertainties, the mass is between about 6.1 and 6.7 $M_\odot$, or $M/M_\odot = 6.4 \pm 0.3$. If the error oval of the star is a little above or below the pair of isochrones for the age range of the cluster, the mass uncertainty is increased from about 5% to about 7%.

As noted in Sect. 3.1, there are several clusters for which we have determined a more precise age by forcing the cluster isochrone to pass through the error oval of an Ap cluster member near the TAMS. The method is very similar to that illustrated by Kochukhov & Bagnulo (2006) in their Fig. 3. In these cases, we have included in the estimate of cluster age uncertainty a contribution from the uncertainty of the appropriate Z value to use for the isochrones. Even including this effect, in a few cases uncertainties in $\log t$ as small as $\pm 0.05 - 0.08$ dex are obtained.
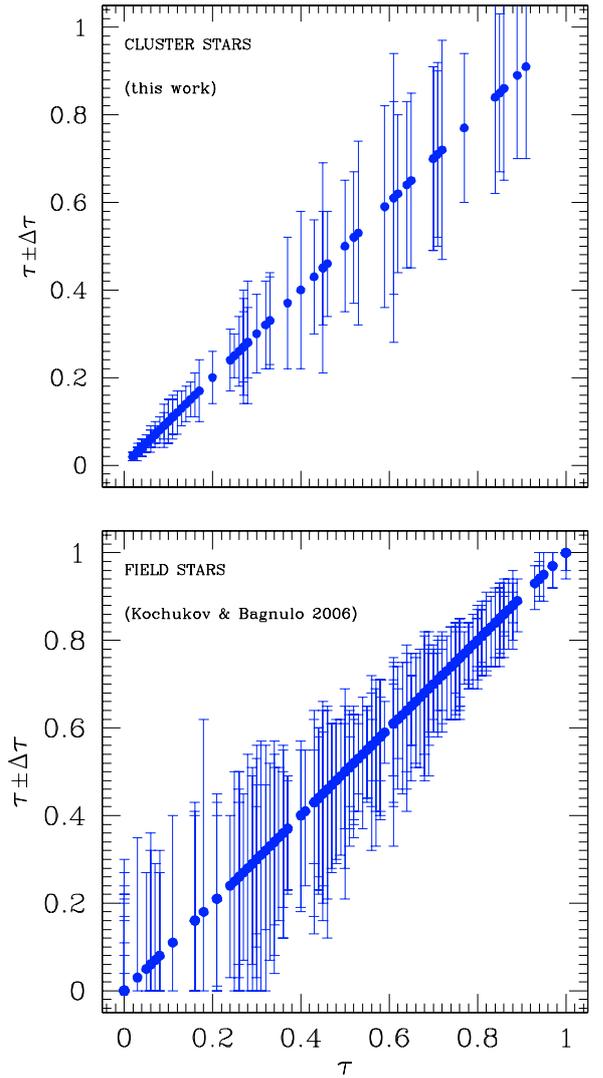
Fig. 3. *Upper panel*: distributions of the 1σ confidence interval of fractional age τ for our sample of open cluster stars. For comparison, we show also the same kind of distribution for the sample of field stars analysed by Kochukhov & Bagnulo (2006). Note that Kochukhov & Bagnulo assume considerably smaller errors in $T_{\rm eff}$ than we do, leading to smaller uncertainties in τ, as discussed in the text.

We then use the mass and its uncertainty to derive the expected main sequence lifetime of each star, again using the $Z = 0.02$ models. Dividing the adopted cluster age by this lifetime, we obtain an estimate of the fraction of the main sequence lifetime already elapsed (the fractional age τ) for each star.

Several sources of error contribute to the uncertainty in this quantity. Mass uncertainty of about 5% leads to main sequence lifetime uncertainty of about 12% for evolution tracks for a given Z. The uncertainty in the actual value of Z for each cluster introduces an additional lifetime uncertainty of up to about 7%, leading to a total uncertainty in main sequence lifetime of about $\pm 15\%$. Uncertainty in the ages of the clusters and associations has been estimated in most cases to be approximately $\pm 0.1 - 0.2$ dex, or about $\pm 25 - 50\%$. In the clusters with age uncertainties in this range, cluster age uncertainty dominates the total error of τ. In those clusters with one or more stars near the TAMS, for which we can obtain more precise age estimates, the contributions from cluster age uncertainty and from main sequence lifetime uncertainty can be comparable.

We treat the uncertainties in the stellar lifetimes and the cluster ages as independent. For stars which are young relative to their main sequence lifetimes (τ less than about 0.3) the accuracy of the fractional age is substantially better than is possible at present by placing field stars in the HR diagram, even for clusters with rather poorly known ages. For stars of fractional age below about 0.1, the improvement compared to the accuracy that can be obtained for field stars is an order of magnitude or more. For τ larger than about 0.5 (these are the situations in which significant improvement in cluster age precision, by isochrone fitting of individual evolved stars, is possible), the uncertainty in fractional age of a cluster Ap (for given error bars on T_{eff} and $L/L_\odot$, and comparable treatment of the Z uncertainty) is similar to that for nearby field stars, since effectively in this circumstance the cluster age is most precisely determined in the same way that the age of a field star is determined.

The age confidence limits which characterise our sample are shown in Fig. 3, together with those obtained in the study based on field stars by Kochukhov & Bagnulo (2006). It is important to keep in mind when comparing these two distributions two substantial differences between their analysis and ours. First, they have assumed uncertainties for T_{eff} that are typically about 0.012 to 0.014 dex, compared to our uncertainty of 0.020 dex (they do not include possible effects due to the overall uncertainty of the T_{eff} scale for Ap stars, which we consider an important probable source of uncertainty). Increasing their T_{eff} uncertainty by 50% would increase the age uncertainty for evolved stars by a similar amount. Secondly, they have not included any contribution from uncertainty in the bulk Z values of the stars of their sample (and thus uncertainty in which evolution tracks to use for age determination), an effect which we have found in Paper I to make an important contribution to the error budget of ages for field stars. If Kochukhov & Bagnulo had used the same uncertainty estimates as we have, their 1σ confidence limits would be very roughly a factor of two larger than those shown in Fig. 3, for all values of τ.

In spite of these important differences, it is easily seen that the precision of our fractional ages is considerably better than that of Kochukhov & Bagnulo for fractional ages below about 0.5. For fractional ages below about 0.2, the improvement is about an order of magnitude, or more. Above $\tau \approx 0.5$, the fact that our fractional ages are less precise than theirs is largely due to the two differences in analysis noted above, since stellar main sequence lifetimes and cluster ages for the most evolved stars are determined in the same way as field star lifetimes and ages.

The stars of the final subsample are listed in Table 3, together with the data derived for each star as discussed above. The first three columns of this table characterise the clusters whose stars are in our subsample, giving the cluster or association name, our adopted (logarithmic) cluster age $\log t$ with uncertainty, and the true distance modulus DM of the cluster. The remaining columns list the stars which we consider to be members with $\log T_{\mathrm{eff}}$, $\log(L/L_\odot)$, $M/M_\odot$ and uncertainty, fractional age τ and uncertainty, and the rms estimate of the longitudinal field B_{rms}.

5. Characteristics of our sample

It is of interest to note some distinctive properties of the data set of Table 3 concerning age and mass distributions.

5.1. Absolute and fractional age distribution

Because open clusters and especially associations tend to be disrupted by galactic tidal acceleration on a time scale short compared to the main sequence lifetime of an A star, we have a predominance of relatively young stars (fractional ages less than 0.5) in our survey, although the sample does include a significant number of more evolved main sequence stars (especially at higher masses where the lifetimes are shorter). In fact, about half the stars in the sample have fractional ages of less than 0.20. However, our data set is fairly uniformly distributed in (log) absolute age between $\log t \approx 6.5$ and 8.5. The distribution of absolute ages and of fractional ages in our sample are shown in the first and second boxes on the left of Fig. 4, respectively.

5.2. Mass distribution

Because the distribution of masses in a cluster or association follows the initial mass function (IMF), at least up to the cluster turn-off, rather than the field mass function, which is more heavily weighted towards low-mass stars of long main sequence lifetimes, our sample has a relatively large number of hot, high-mass magnetic Bp stars. The median mass in Table 3 is about 3.5 $M_\odot$, and the median value of T_{eff} is about 12 700 K. These are stars that are not particularly common among the Ap-Bp stars near the Sun, and may well prove to be rather interesting objects to study individually. The distribution in mass of stars in our subsample is shown in the third box in Fig. 4. Notice that we have only three stars (CPD -32 13119, CPD -20 1640, and HD 66318) that have $M/M_\odot \leq 2.1$, and of these three, only HD 66318 has a detected (but huge) magnetic field (cf. Bagnulo et al. 2003). Significant numbers of Ap stars with detected magnetic fields begin to be found in our sample only above about 2.3 $M_\odot$.

5.3. Magnetic field distribution

The distribution of magnetic field strengths found for probable cluster Ap stars in our survey (including Ap stars in which no field has yet been detected) is shown in the right box in Fig. 4. This distribution is qualitatively similar to what is found for samples of field Ap stars (e.g. Bohlender & Landstreet 1990), but scaled up to somewhat larger fields than usual. The bulk of the stars (about 2/3 of the present sample) have rms fields of less than 1 kG, and the median field is about 450 G. This may be compared to the distribution of Bohlender & Landstreet; for a sample of 12 bright Ap field stars (also including some in which fields had not yet been detected) they found none with B_{rms} above 1 kG, and a median field of about 300 G. The largest difference between this sample and that of Bohlender & Landstreet is the presence of a substantial tail of high-field stars; almost a quarter of the present sample have rms fields of 2 kG or more. The differences between these two samples is probably due to the much larger fraction of young Ap stars in the present sample (see the discussion below).

6. Discussion

We finally consider what general conclusions may be drawn from our data about the evolution of magnetic fields up to and through the main sequence phase of stellar evolution. With a sample of more than 80 stars which are probable or definite cluster or association members, probable or definite Ap-Bp stars, and for which at least one magnetic measurement with σ of the order of 10^2 G is available, we are now in a very good position to carry out a preliminary statistical assessment of the evolution of Ap-Bp star magnetic fields based on well-determined stellar ages.
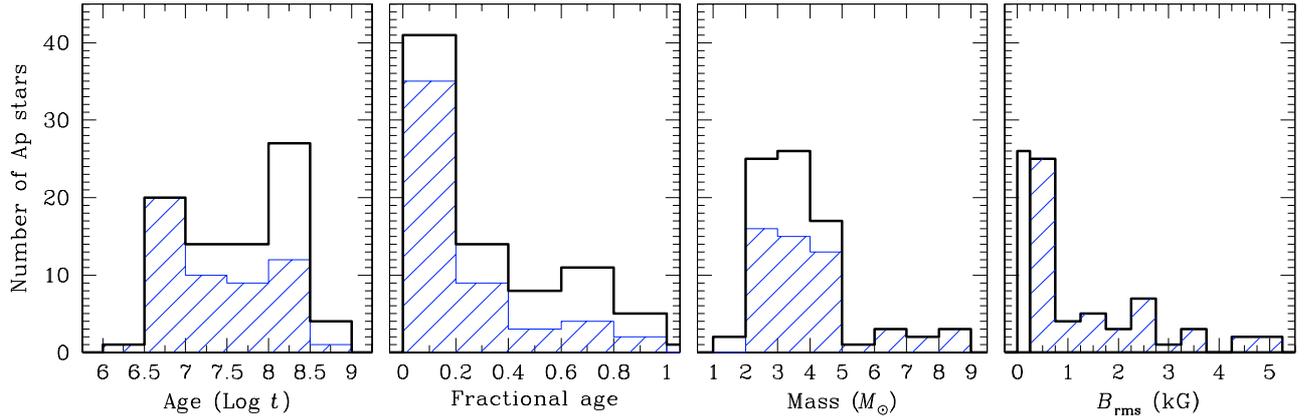
Fig. 4. Distributions in our sample of age, fractional age, mass, and B_{rms}. In each box, the shaded histogram shows the distribution of those Ap stars of Table 3 for which a $B_{rms} \geq 0.25$ kG has been measured. The thicker lines show the distribution of the entire sample.

The fundamental results of our survey are displayed in the eight panels of Fig. 5, where we plot, using logarithmic scales on all axes, the field strength B_{rms} as a function of stellar age (left column of the figure), and as a function of stellar mass (right column). In both columns, the top panel shows all stars of the sample (except θ^1 Ori C), while below we see the same plot for three separate mass ranges (left column) or age ranges (right column), with parameter ranges shown in each panel.

The use of logarithmic scales would introduce a significant bias into fits to these data, in that an rms field which happens to have a value of only a few G is substantially lower than the point for a star with an rms field of, say 150 G, although with the uncertainties of our data, there is no statistical difference between the two values. In order to avoid having small rms field values (essentially at present, non-detections) have an undue influence on fits, we have artificially offset all B_{rms} values by +200 G.

6.1. Does magnetic fields strength change with with time?

The fundamental effect we are looking for with this study is to discover whether typical rms fields of magnetic Ap stars change with stellar age, and if so, how they change. The top left panel of Fig. 5 appears to provide a first tentative answer to this question. The variation of log B_{rms} as a function of log t has been fit with a linear function. This fit to the (modified) data, shown as a line through the data points, was found to have a slope which is non-zero (negative) at about the 5σ level (see Table 1). It appears that the typical field strength decreases markedly with stellar age, from a value of more than 1000 G for stars having ages of about 3×10^6 yr to a value of the order of 200 G at 6×10^8 yr.

However, from Table 3, it is clear that most of the stars with very young ages (say, less than 3×10^7 yr) are relatively massive (masses around 4 or 5 $M_\odot$), while the stars with ages of more than 10^8 have smaller masses (below about 4 $M_\odot$). Another symptom of this systematic change in mass with age is seen in the variation of radii (in fact, we have plotted the value of $(R/R_\odot)^2$) of the stars: the radius of a single star increases with age, but the radii of the sample show no significant variation with age, because the increase in radius of individual stars is essentially compensated by the fact that older stars have smaller masses and radii.

Thus we have considered three smaller sub-samples having limited mass ranges: $M < 3 M_\odot$ (shown in the second panel from the top on the left), $3 M_\odot \leq M \leq 4 M_\odot$ (third panel from the top), and $M > 4 M_\odot$ (left bottom panel). Note that the subsample of

stars defined by $M < 3 M_\odot$ is in fact mainly populated with stars between 2.3 and 3 $M_\odot$, and the subsample defined by $M > 4 M_\odot$ in fact contains mainly stars between 4 and 5 $M_\odot$. The selected mass ranges are narrow enough that the inhomogeneity of mass is not a major factor.

It is seen that for the two samples with larger masses ($3-4 M_\odot$ and $>4 M_\odot$) the B_{rms} decreases with time with a slope that is non-zero at the 3σ level. For the $3-4 M_\odot$ sample, more than half of the stars younger than 4×10^7 yr have rms fields greater than 1 kG, while all but one of the stars in this range older than this age have fields below 1 kG. Similarly, in the most massive sample, a majority of the stars with ages below 3×10^7 yr have fields above 1 kG while none of the older stars (with a range of ages up to more than 10^8 yr) have fields above 1 kG.

In contrast, the lowest mass sample (mainly populated with stars between 2.2 and 3 $M_\odot$) does not show a significant decrease of rms field with age, and in fact there are about as many fields above 1 kG among the older (age $> 10^8$ yr) stars as among the younger stars of this mass range. Note that this group has no representatives among the youngest clusters; we have not found any magnetic Ap stars with $M < 3 M_\odot$ in stellar groups of ages less than 10^7 yr, although the Sco OB2 association is close enough for such stars to be readily identified.

One possible simple scenario for the observed decline of B_{rms} with age among stars of $M > 3 M_\odot$ would be that such stars (approximately) conserve magnetic flux as they evolve through the main sequence phase. During the evolution from ZAMS to TAMS, the radius of a star increases by a factor of roughly 3.

We have tested our samples to see if flux $\Phi_B \sim B_{rms}(R/R_\odot)^2$ is statistically conserved as we look at stars of various ages. In the various panels of Fig. 5, we plot both $(R/R_\odot)^2$ and $B_{rms}R^2$. In the left column of panels, the individual mass bands all show increasing values of $(R/R_\odot)^2$ with increasing age at the 3σ level, although the rise is not as great as expected, probably due to the increasing contribution of lower mass stars in each sample as we go to greater ages. In the two panels at the lower left, where we find strongly significant decline in B_{rms}, the value of flux $B_{rms}R^2$ appears to decline somewhat, but the decline is only present at about the 2σ level. We may regard this as consistent either with some real decline in magnetic flux, or with conservation of magnetic flux.

To summarise the results of this analysis, we find that magnetic Ap stars with masses above 3 $M_\odot$ have fields which decline substantially over an age of the order of $2-3 \times 10^7$ yr. For these
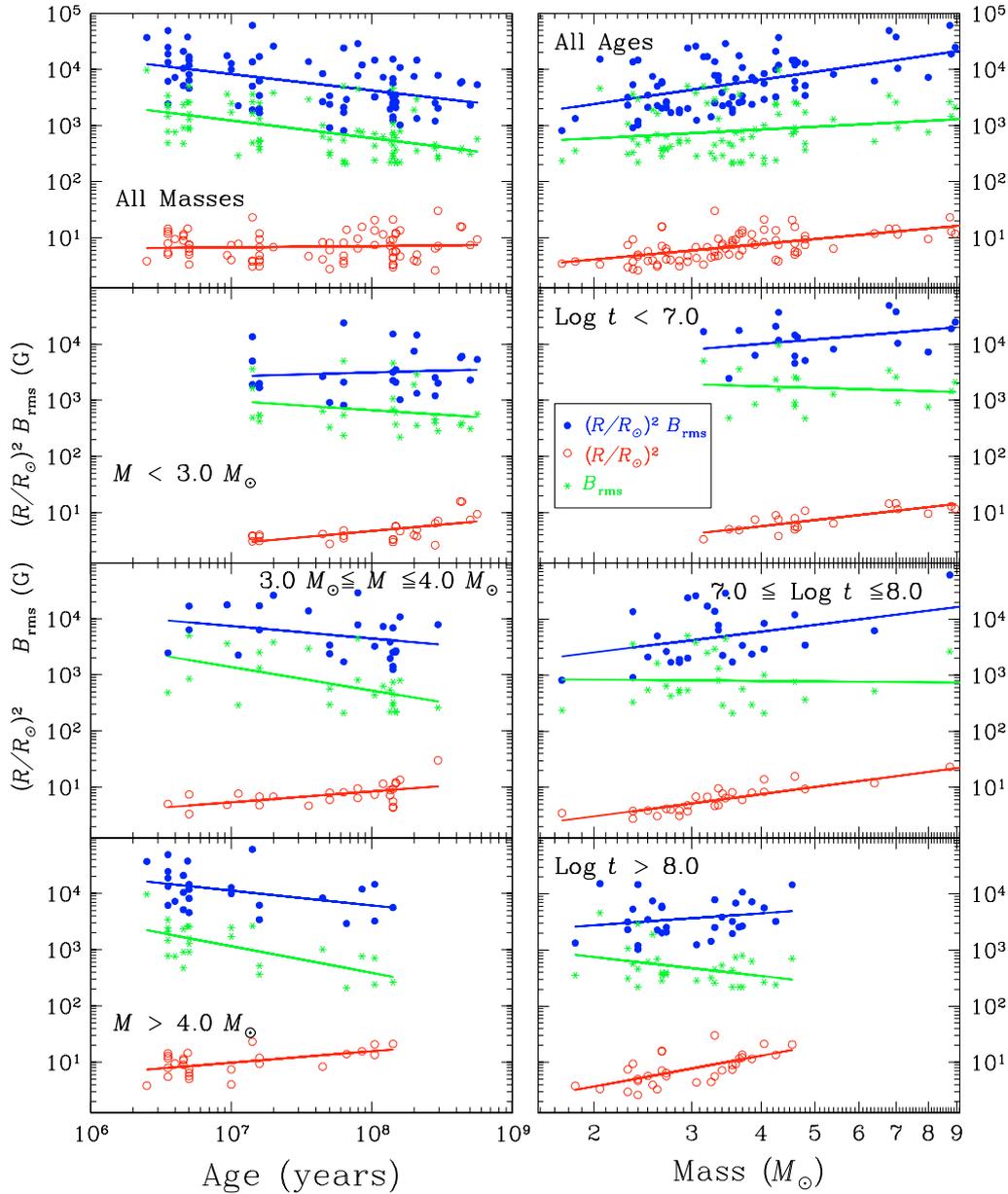
Fig. 5. Plots of B_{rms} (stars), $(R/R_\odot)^2$ (open circles), and $B_{\mathrm{rms}}(R/R_\odot)^2$ (filled circles) are shown as functions of stellar age (*left column*) and mass (*right column*) for the subsample of Table 3. In each column the *top panel* shows all stars in the sample (except for θ^1 Ori C). *Lower panels* show various cuts through the subsample as indicated on each panel. As discussed in the text, 200 G has been added to each value of B_{rms}, in order to avoid having correlations made artificially steeper by very small values of B_{rms}.

stars, the magnetic flux may decline with time, or remain roughly constant. The rms fields of less massive stars change little over time-scales approaching 10^9 yr, and there is no strong evidence that the magnetic fluxes change significantly over this time scale.

The evolution with time of fields in main sequence Ap stars has also been discussed by Kochukhov & Bagnulo (2006). They find that the value of B_{rms} declines with increasing age for all their subsamples of Ap stars except for those with $M/M_\odot < 2$, in contrast to our result that B_{rms} decreases with age except for stars of $M/M_\odot < 3$. Thus the two studies agree that there is a significant decrease of B_{rms} with time among Ap stars at least above 3 $M_\odot$, although they disagree on precisely where this behaviour changes over to more nearly constant fields.

However, Kochukhov & Bagnulo (2006) find a significant increase in magnetic flux with age for stars of $M/M_\odot < 3$, while we find, if anything, a modest decrease of flux with time. At this

point, it is probably fair to say that approximate magnetic flux conservation during the main sequence phase is not ruled out. It will be very useful to obtain more precise values of B_{rms} for stars in the cluster sample (by getting multiple observations of each star, and expanding the sample of stars measured) to resolve the differences between these two studies.

6.1.1. Do magnetic fields appear late in main sequence evolution?

So far we have discussed a possible decline of the magnetic field, but we have not commented on whether the magnetic field is present at the very moment when a star reaches the ZAMS, or whether the field appears at the stellar surface at some later stage in its main sequence evolution. This problem has been previously discussed, among others, by Hubrig et al. (2000),

Table 1. Slopes of regression lines in Fig. 5.

Sample	y-axis	x-axis	Slope
	$\log(R^2 B_{\mathrm{rms}})$	$\log t$	-0.29 ± 0.07
All masses	$\log R^2$	$\log t$	0.02 ± 0.04
	$\log B_{\mathrm{rms}}$	$\log t$	-0.31 ± 0.06
	$\log(R^2 B_{\mathrm{rms}})$	$\log t$	0.07 ± 0.15
$M < 3.0\, M_\odot$	$\log R^2$	$\log t$	0.23 ± 0.06
	$\log B_{\mathrm{rms}}$	$\log t$	-0.16 ± 0.14
	$\log(R^2 B_{\mathrm{rms}})$	$\log t$	-0.22 ± 0.14
$3.0\, M_\odot \leq M \leq 4.0\, M_\odot$	$\log R^2$	$\log t$	0.19 ± 0.06
	$\log B_{\mathrm{rms}}$	$\log t$	-0.42 ± 0.14
	$\log(R^2 B_{\mathrm{rms}})$	$\log t$	-0.27 ± 0.12
$M > 4.0\, M_\odot$	$\log R^2$	$\log t$	0.21 ± 0.07
	$\log B_{\mathrm{rms}}$	$\log t$	-0.47 ± 0.11
	$\log(R^2 B_{\mathrm{rms}})$	$\log(M/M_\odot)$	1.44 ± 0.26
All Ages	$\log R^2$	$\log(M/M_\odot)$	0.92 ± 0.14
	$\log B_{\mathrm{rms}}$	$\log(M/M_\odot)$	0.52 ± 0.29
	$\log(R^2 B_{\mathrm{rms}})$	$\log(M/M_\odot)$	0.84 ± 0.57
$\log t < 7.0$	$\log R^2$	$\log(M/M_\odot)$	1.12 ± 0.21
	$\log B_{\mathrm{rms}}$	$\log(M/M_\odot)$	-0.29 ± 0.61
	$\log(R^2 B_{\mathrm{rms}})$	$\log(M/M_\odot)$	1.23 ± 0.52
$7.0 \leq \log t \leq 8.0$	$\log R^2$	$\log(M/M_\odot)$	1.32 ± 0.14
	$\log B_{\mathrm{rms}}$	$\log(M/M_\odot)$	-0.09 ± 0.51
	$\log(R^2 B_{\mathrm{rms}})$	$\log(M/M_\odot)$	0.70 ± 0.61
$\log t > 8.0$	$\log R^2$	$\log(M/M_\odot)$	1.83 ± 0.38
	$\log B_{\mathrm{rms}}$	$\log(M/M_\odot)$	-1.13 ± 0.57

Fig. 6. The position of the stars with $M \leq 3\, M_\odot$ in the diagram of B_{rms} vs. fractional age.
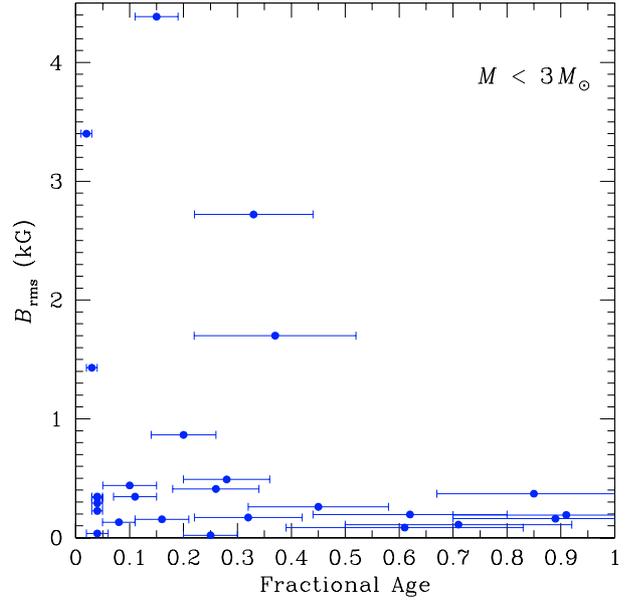
Bagnulo et al (2003), Pöhnl et al. (2005), and Kochukhov & Bagnulo (2006).

Based on the analysis of field stars, Hubrig et al. (2000) have proposed that, in stars with $M \lesssim 3\, M_\odot$, a magnetic field appears at the surface of an Ap stars only after about 30% of its main sequence lifetime has elapsed. This conclusion was later reaffirmed by Hubrig et al. (2005a,b). However, this result was contradicted by the discovery of a field in the young 2.1 $M_\odot$ star HD 66318 (Bagnulo et al. 2003). Furthermore, both Pöhnl et al. (2005) and Kochukhov & Bagnulo (2006) were unable to reproduce the results of Hubrig et al. of late emergence of surface fields, but both studies found a remarkable shortage of young magnetic stars with $M \lesssim 2\, M_\odot$.

In the present sample we have a total of six stars with masses definitely below 3 $M_\odot$, fractional ages below 0.30, and unambiguously detected magnetic fields. These stars are HD 65712 (Paper I), HD 66318 (Bagnulo et al. 2003), HD 92385 (Paper I), HD 112381 (Bohlender et al. 1993), HD 119419 (Thompson et al. 1987), and HD 318100 (Paper I). One additional star, HD 74169 (Paper I), satisfies all criteria but only has one measurement with field detection at the few sigma level, and still ought to be confirmed. Figure 6 shows the position of all of the stars of our sample with $M \leq 3\, M_\odot$ in the diagram of B_{rms} vs. fractional age. Our study thus definitely contradicts that of Hubrig et al. concerning late field emergence.

We have investigated the reasons for the discrepancies between our conclusions and those of Hubrig et al. (2000) and we have identified several points that may have led them to obtain incorrect results.

i) Hubrig et al. (2000) have used the bolometric corrections of Schmidt-Kaler, which are too large even compared to the most accurate values for normal stars, and lead to an overestimate of bolometric magnitude relative to the BC's adopted here (see Fig. 1) by about 0.25 mag (0.1 dex in $\log(L/L_\odot)$). Using more appropriate BC's would move their observed stars downwards, towards the ZAMS, and younger ages.

ii) As discussed above, the whole effective temperature scale for magnetic Ap stars may be systematically a few percent low. This possibility has little effect on ages derived for cluster stars, but, if correct, would move the field stars towards the ZAMS as well.

iii) Hubrig et al. (2000) applied the Lutz-Kelker correction to their data, moving the stars upwards in the HR diagram, away from the ZAMS. As discussed by Stępień (2004) and Kochukhov & Bagnulo (2006), this correction should not have been used. Removing it also moves the Hubrig et al. sample closer to the ZAMS.

iv) Finally, in Hubrig et al. (2000) there are a two significant errors: the T_{eff} value of HD 119419 is about 15% too low, and that of HD 147010 is more than 50% too small. Correcting these errors would move HD 147010 to a mass slightly above 3 $M_\odot$ but to a position close to the ZAMS. HD 119419 would have remained well below 3 $M_\odot$ and moved close to the ZAMS, providing Hubrig et al. with a counter-example to their hypothesis in their own data.

The changes to the data of Hubrig et al. (2000) resulting from the points discussed above appear to be enough to make the effect they report disappear from their data. This result, together with the results of Bagnulo et al. (2003), Pöhnl et al. (2005), Kochukhov & Bagnulo (2006), and our own survey, lead us to reject the hypothesis that stars with $M \leq 3\, M_\odot$ becomes magnetic only after they have spent a significant fraction of their life in the main sequence.

In our survey we note a relatively small, but non-zero, number of stars of mass below about 2 $M_\odot$ ($T_{\mathrm{eff}} < 8500$ K). Only three of the 81 stars in Table 3 have masses in the range $1.7 \leq M/M_\odot \leq 2.1$. To some extent this may reflect the absolute rarity of such stars; in the field, the Ap stars near F0 make up only about 1% of all stars, compared to about 10% near A0 (Wolff 1968). Still, a fairly large number of such stars are now

known in the field, both as a result of searches for roAp stars (Kurtz & Martinez 1993) and from searches for magnetic stars showing visible Zeeman splitting, which seems to be commoner among low-mass Ap stars (Hubrig et al. 2000). A number of such low-mass Ap stars are also found in the survey of Kochukhov & Bagnulo (2006); the rarity of low-mass Ap stars in our sample seems to be anomalous. It may result primarily from selection effects. Such stars will be among the faintest Ap stars in most of the clusters we have studied, and may simply not have been selected as Ap stars in identification studies, either because they were near the threshold of the study, or because the method used (such as the Δa method) is not very sensitive to the coolest Ap stars.

Note that, in contrast to the results of Pöhnl et al. (2005) and Kochukhov & Bagnulo (2006), both of whom found a deficiency of *young* Ap stars with masses below 2 $M_\odot$, all three of the low-mass stars in our survey are fairly close to the ZAMS; the largest fractional age among these stars is 0.16. This effect probably results from the fact that the main sequence lifetime of such stars (10^9 yr or more) is considerably longer than the ages of most clusters in our sample. Another distinctive feature of our sample of particularly low-mass Ap stars is that no field has so far been definitely detected in either of the two stars with masses below 2 $M_\odot$. The lowest mass cluster Ap star for which a field has been definitely detected is HD 66318 (Bagnulo et al. 2003). This is quite different from the situation in the field, where a number of Ap stars with detected fields have masses that extend down to about 1.5 $M_\odot$ (cf. Kochukhov & Bagnulo).

From the present study (in particular from the fractional ages reported in Table 3) it is clear that magnetic fields are present at essentially the zero age main sequence phase for all stellar masses between about 2 and 5 $M_\odot$; stars with fractional ages below about 0.05 occur throughout this range. It seems very reasonable to suppose that these fields are already present late in the pre-main sequence phase, and this has recently been confirmed by the discovery of ordered global fields in several pre-main sequence Herbig AeBe stars (Donati et al. 1997; Wade et al. 2005; Catala et al. 2006).

6.2. Do magnetic fields depend on mass?

A further issue that we may address with our data is the question of whether there is a variation with mass of typical rms field strength, at a given stellar age. Thompson et al. (1987) have argued that they detected such an effect, with stars more massive than roughly 6 or 7 $M_\odot$ having fields about a factor of two larger than lower mass stars. Kochukhov & Bagnulo (2006) find that the field strength $B_{\rm rms}$ shows no clear variation with mass, but argue that the magnetic flux definitely increases with mass.

A plot of the rms fields of all the stars in our sample as a function of stellar mass (top right panel in Fig. 5) appears to show a trend of increasing field strength with mass. However, again we must recall that this sample has considerable change in typical mass as a function of age. To reduce the influence of this effect, the whole sample is subdivided into three samples of more limited age range: $\log t < 7.0$, $7.0 \leq \log t \leq 8.0$, and $\log t > 8.0$. As noted above, the youngest sample contains only stars more massive than about 3 $M_\odot$, and the oldest sample is restricted to stars less massive than 4 $M_\odot$. None of these smaller samples of restricted age shows a clearly significant trend of rms field with mass. In particular, we have six stars in the youngest sample with masses above the value at which Thompson et al. (1987) thought they had found a jump in field strength, but we see no significant change at this mass.

From our data, it is not clear if magnetic flux is typically a function of stellar mass. The age range between $\log t = 7$ and 8 appears to show a significant increase of magnetic flux with mass, but this does not seem to be present (over more restricted mass ranges) for younger or older samples.

7. Conclusions

We have assembled a sample of nearly 160 stars for which magnetic field measurements are available (primarily from Paper I) and which are *possible* members of clusters or associations, and *possible* magnetic Ap or Bp stars. After detailed examination of astrometric, photometric, and spectroscopic data available for these stars, 81 remain as *probable* cluster members and *probable* magnetic Ap stars. These stars are listed in Table 3.

For these stars, we have obtained cluster/association distances and ages (typically accurate to $\pm 0.05 - 0.2$ dex) from the literature and from our own isochrone fits to stars near the TAMS. We have determined effective temperatures (typically accurate to ± 0.02 dex) for almost all of these stars from both Strömgren and Geneva photometry. We have derived new bolometric corrections for Ap stars, and obtained luminosities of individual stars with typical uncertainties of ± 0.1 dex.

By placing the stars of this subsample in the HR diagrams of their clusters, we have determined their masses with an accuracy of about $\pm 5\%$, and improved the precision of the ages of several clusters. The mass values allow us to obtain main sequence lifetimes from evolutionary calculations, with an overall uncertainty (including the uncertainty in bulk chemical composition) of about $\pm 15\%$. From main sequence lifetimes and cluster ages, we deduce fractional ages τ (the fraction of the main sequence lifetime of each star already elapsed). The uncertainty in the fractional age is almost always less than half of the actual fractional age, and thus for stars near the beginning of their main sequence lives, the fractional ages are *much* more accurate than ages than can be derived for field stars by placing them in the HR diagram (see the penultimate columns of Table 3).

From the stellar characteristics derived here we have assembled for the first time a substantial sample of magnetic Ap stars for which the absolute and fractional ages are known with really useful precision. These accurate masses and fractional ages allow us to definitively test the hypothesis of Hubrig et al. (2000) that in Ap stars of masses $M < 3\ M_\odot$, the magnetic fields first become visible at the surface after about 30% of the main sequence lifetime has elapsed. We have identified six clear counter-examples to this hypothesis in our sample. As a result, we are able to definitively rule out the hypothesis of late field emergence in low-mass Ap stars.

Although the field measurements of the stars of our sample are still rather incomplete, we can use the available data to constrain the typical evolution of magnetic field strength with stellar age. We find that fields are definitely present essentially at the ZAMS (fractional ages of 0.05 or less) for Ap stars ranging between 2 and 5 $M_\odot$, and for fractional ages of less than about 0.10 for masses up to 9 $M_\odot$.

The evolution of magnetic fields with time suggested by our data reveals a quite unexpected pattern. For stars of $M > 3\ M_\odot$ we find strong evidence that the field strength declines by a factor of a few on a timescale of about $2-3 \times 10^7$ yr, in agreement with the conclusion of Kochukhov & Bagnulo (2006). It is not clear for these stars whether the total magnetic flux remains approximately constant, causing $B_{\rm rms}$ to decrease as the stellar radius expands, or whether this flux actually declines slowly.

In contrast, for stars of $M < 3\,M_\odot$, there is no convincing evidence in our data of field strength decrease, nor of magnetic flux decrease, even on a time scale of several times 10^8 yr. This conclusion, very different from the result for more massive Ap stars, represents a significant puzzle; perhaps this result reveals some important structural difference between the lower and higher mass Ap stars.

Finally, we are able to examine the question of whether there is a significant variation in typical magnetic field strength as a function of mass within a given age band. We do not find any strong evidence for such a variation, in contrast to Thompson et al. (1987), who argued that typical field strength increases rather abruptly at about $6-7\,M_\odot$.

These results indicate clearly that studies of magnetic stars in clusters can yield much useful information about the nature and evolution of their fields. Further observations, both to detect fields in more cluster Ap stars (thereby increasing the size and diversity of the sample), and to improve the accuracy of the rms fields in stars already detected, are already in progress.

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Table 2. Tests of cluster membership and chemical peculiarity. The first two columns give the cluster or association to which the star may belong, and a common designation of the star. Columns 3−5 indicate whether the measured parallax, proper motions, and position in the HR diagram are consistent with membership, and column 6 gives our conclusion about cluster membership (y = yes, p = probable, ? = questionable, n = no, blank = no data; letters in parentheses refer to references at end of table). Columns 7 through 10 summarize evidence that each star is an Ap star according to its spectral classification, to the value of Δa and/or Z, to the presence of periodic variability, and to whether a magnetic field is actually detected (letters in parentheses in the "mag fld" column refer to references at the end of the table). Column 11 reports our conclusion as to whether the star is indeed an Ap star.

Cluster or Association	Star	Cluster membership				Magnetic Ap star				
		π	μ	phot	member?	Sp	$\Delta a - Z$	var	mag fld	Ap?
Blanco 1	HD 225264	p	y		y (R, B, D, K)	?	n	n	n (F)	?
α Per	HD 19805	y	p	y	y (Z, K)	?	p	n	? (a, r)	?*
	HD 20135		n		n (K)	p	y		n (a)	p
	HD 21699	y	p	y	y (R, Z, K)	y	y	y	y (b)	y
	HD 22401	y	p		p (Z, K)	p	n	n	n (a)	?
Pleiades	HD 23387A		y	p	y (D, K)	p	?		? (a, r)	?*
	HD 23408	y	y		y (D, K)	?	n	n	n (a, s, F)	?
NGC 1662	HD 30598	y	y	y	y (D, K)	p	y	n	n (F)	p
Ori OB1a	HD 35008	p	y	?	? (Br)	p	n	n	y (F)	y
	HD 35298	y	y	p	p (Br)	y	y	y	y (c)	y
	HD 35456	y	y	?	? (Br)	y	y	p	y (c)	y
	HD 35502	y	y	y	p (Br)	p	p	p	y (c)	y
	HD 36429	y	y		p (Br)	p	n	n	? (c)	?
	HD 36549	y	p		p (Br)	y	?	n	n (F)	?
Ori OB1b	HD 36046	p	y		p (Br)	y	n	n	n (F)	?
	HD 36313		y	?	? (Br)	y	y	y	y (c)	y
	HD 36485		p	y	p (Br)	y		y	y (d)	y
	HD 36526		y	y	p (Br)	y	p	y	y (c)	y
	HD 36668	y	y	?	? (Br)	y	y	y	y (c)	y
	HD 290665		y	?	? (Br)	y	y		y (F)	y
	HD 37140		y	y	p (Br)	y	p	y	n (c)	y
	HD 37333	y	y		p (Br)	n	y	n	n (F)	?
	HD 37479		y	y	p (Br)	y		y	y (d)	y
	HD 37633		p	y	p (Br)	y	y		y (F)	y
	HD 37776	y	y	y	p (Br)	y		y	y (g)	y
Ori OB1c	HD 36540	y	y	p	p (Br)	y	n	y	y (F)	y
	HD 36629	p	y		p (Br, T)	p		n	n (F)	?
	HD 36918		y		p (Br, T)	?	n		n (F)	?
	HD 36916	y	p	?	? (Br)	y		y	y (c, F)	y
	HD 36960	y	y		p (Br, T)	p		n	n (F)	?
	HD 37017	y	y	?	? (Br)	y		y	y (d)	y
	HD 37058		y	y	p (Br)	p		y	y (k, F)	y
	HD 37210		y	?	? (Br)	y	y	y	n (c, F)	y
	HD 37470	p	y		p (Br)	?		n	n (c, F)	?
	HD 37642	y	y	p	p (Br)	y	y	y	y (c)	y
Ori OB1d	HD 36982		y		y (Br, T)	p			n (F)	?
	HD 37022		y	y	y (Br, *)	n		y	y (e, F)	y
NGC 2169	NGC 2169 12		p	y	p (K)	p	p	y	y (F)	y
NGC 2232	HD 45583	y	y	y	y (R, B, K)	y	y	y	y (F)	y
NGC 2244	NGC 2244 334		y	p	y (M, S)	y	y		y (F)	y
NGC 2287	HD 49023	p	y		y (B, D, K)	p		n	n (F)	?
	CpD-20 1640		p	y	p (D, K, *)	p	n		n (F)	p
	HD 49299		p	y	p (D, K)	y	y		y (F)	y
	HD 49333	n	n		n (B, D, K)	y		y	y (h, k)	y
CMa OB2	HD 51088		p	p	p (D, K)	y	y		? (F)	y
NGC 2323	HD 52965		p	y	p (B, D, K)	p		n	n (F)	p
Cr 132	HD 56343	p	y	p	p (B, *)	n	y	n	y (F)	y
Cr 135	HD 58260	n	n		n (B, D, K)	y			y (d)	y
NGC 2422	BD-14 2015		p		p (K)	p	n	n	n (F)	?
	HD 61045	y	y	y	y (R, B, D, K)	y	y	n	y (F)	y
	BD-14 2040		y		y (K)	p	n		n (F)	n
	BD-14 2028				p	y	n	n	n (F)	?
NGC 2451A	HD 62376	y	n		n (B, K, *)	n	y	y	n (F)	p
	HD 63079	y	y		y (B, K)	p	n	n	n (F)	?
	HD 63401	y	p	y	y (R, B, K)	p	y	y	y (s, F)	y
NGC 2451B	CD-37 3845		p		p (C)	p	?		n (F)	?
	HD 62992	p	n		n (C, K)	y	y		? (F)	y
NGC 2489	NGC 2489 58				?	?	y		n (F)	?
	NGC 2489 40			y	?	?	y		n (F)	?
NGC 2516	HD 65712		y	y	y (K, *)	y	y		y (F)	y

Table 2. continued.

Cluster or Association	Star	Cluster membership				Magnetic Ap star				
		π	μ	phot	member?	Sp	Δa, Z	var	mag fld	Ap?
	CpD-60 944A		y	y	y (D, K)	y	y		p (F)	y
	CpD-60 944B		p	p	p (D, K, *)	p	y		n (F)	y
	HD 65949	p	p		p (D, K, *)	n	n		n (F)	n
	CpD-60 978		p	y	p (D, K)	y	y	y	n (F)	y
	HD 65987		y	y	y (D, K)	y	y	p	y (F)	y
	CpD-60 981				? (K)	?	?		n (F)	?
	HD 66295		y	y	y (D, K)	y	y	y	y (F)	y
	HD 66318		p	y	p (D, K)	y	y	n	y (F)	y
NGC 2546	NGC 2546 258				?	?	y		n (F)	p
	NGC 2546 201				p (K)				y (F)	y
	NGC 2546 197		y	y	y (D, K)	p	y		n (F)	p
	HD 69004		p	?	? (D, K)	y	y		n (F)	p
	HD 69067		n		n (D, K)	y	y		y (F)	y
IC 2391	HD 73340	y	n		n (R)	y	y	y	y (h)	y
	HD 74169		y	y	y (K)	y	y	y	p (F)	y
	HD 74168	n	n		n (R, K)	y	y		n (s, F)	p
	HD 74195	y	y		y (R, K, *)	n	n		n (F)	n
	HD 74196	y	y		y (R, K)	?	n		n (s, F)	n
	HD 74535	y	p	y	p (R, K)	p	y	y	n (F)	y
	HD 74560	y	y		y (R, K)	?	n		n (F)	?
Tr 10	HD 75239		y		y (D, K)	p	n		n (F)	?
NGC 2925	HD 83002	y	y	p	y (D, K)	p			n (F)	p
NGC 3114	HD 87241	y	p		y (B, D, K)	p	n	n	n (F)	?
	HD 87240		n		n (D, K)	y	y		y (F)	y
	HD 87266	p	p		p (B, D, K)	?	n	n	n (F)	?
	HD 304841		p	y	p (D, K)		y		y (F)	y
	HD 304842		p	y	p (D, K)	?	y	y	n (F)	p
	HD 87405	y	?		? (B, D, K)	y	y	n	n (F)	y
NGC 3228	HD 89856	?	n		n (B, D, K)	?	y	?	p (F)	p
	HD 298053		p		p (D, K)	n	?		n (F)	n
vdB-Hagen 99	HD 92190		p	y	p (D, K)	p			n (F)	p
IC 2602	HD 92385	y	y	y	y (R, D, K)	?	y	y	y (s, F)	y
	HD 92664	y	p	y	y (R, K)	y	y	y	y (h)	y
	HD 93030	p	?		? (K)	?			n (F)	?
Cr 228	Cr 228 30		p	y	p (K)	n			y (F)	y
	HD 305451		p		p (K)	p			n (F)	p
NGC 3532	HD 96040		p	p	p (D, K)		y		y (F)	y
	HD 96729		p	?	? (K)	p	y		y (F)	y
	HD 303821		p	?	?		y		n (F)	p
Coma Ber	HD 108662	y	n		n (R, K)	y	y	y	y (i)	y
	HD 108945	p	y	y	y (R, K)	y	y	y	y (j, s, F)	y
NGC 5460	HD 122983		p	y	p (D, K)	n	y		p (F)	p
	HD 123183		p	y	p (D, K)	n	n		p (F)	p
	HD 123225		y	y	y (D, K)	?	p		n (F)	p
NGC 5662	CpD-56 6330		n		n (D, K)	p	y		n (F)	p
	HD 127866	y	y		y (B, D, K)	n	n	n	n (F)	n
	HD 127924	y	y	y	y (B, D, K)	?	y	n	n (F)	p
Lower Cen-Cru (Sco OB2)	HD 103192	y	n		n (Z)	y	y	y	y (h)	y
	HD 112381	p	y	y	y (Z)	y	y	y	y (h)	y
	HD 114365	y	y	y	y (Z)	y			n (n, s, F)	p
	HD 119419	y	y	y	y (Z)	y	y	y	y (h)	y
Upper Cen-Lup (Sco OB2)	HD 122532	p	n		n (Z)	y	y	y	y (n)	y
	HD 125823	y	y	y	y (Z)	y		y	y (k)	y
	HD 126759	y	y	y	y (Z)	p			? (n)	p
	HD 128775	p	y	y	y (Z)	y		y	y (n, s)	y
	HD 128974	?	n		n (Z)	p			n (n, F)	?
	HD 130559	n	n		n (Z)	y		n	y (q)	y
	HD 131120		y		y (Z)	?			n (s, F)	?
	HD 133652	p	y	y	y (Z)	y	y	y	y (h)	y
	HD 133880	y	y	y	y (Z)	y	n	y	y (l)	y
	HD 136347	y	y	y	y (Z, *)	p			n (n)	p
	HD 137193	n	n		n (Z)	y	y	y	y (n)	y
	HD 139525	n	n		n (Z)	p			n (n)	p
	HD 143699	y	y	y	y (Z)	?	n	n	n (k)	?
Upper Sco (Sco OB2)	HD 142301	y	y	y	y (Z)	y		y	y (m)	y
	HD 142990	y	y	y	y (Z)	y		y	y (h)	y

Table 2. continued.

Cluster or Association	Star	Cluster membership				Magnetic Ap star				
		π	μ	phot	member?	Sp	Δa, Z	var	mag fld	Ap?
	HD 144334	y	p	y	y (Z)	y		y	y (k)	y
	HD 144661	p	y	p	y (Z)	y	y	n	n (k)	y
	HD 145102	y	n		n (Z)	p			n (n, s, F)	?
	HD 145501	y	y	y	y (Z, *)	y	y	?	y (k)	y
	HD 146001	y	y	y	y (Z)	y	p	n	y (k)	y
	HD 147010	y	y	y	y (Z)	y	y	y	y (n)	y
	HD 147890	n	n		n (Z)	p		y	n (n)	y
	HD 148199	y	n		n (Z)	y	y	?	y (d, h)	y
	HD 151525	y	n		n (Z)	y	y	y	p (s, F)	y
NGC 6087	CpD-57 7817			y	p (K)		y		y (F)	y
	HD 146555		y	y	y (D, K)	y	y		? (F)	y
NGC 6178	HD 149257		y		y (K)	p			n (F)	p
	HD 149277	y	y	p	y (K)	n		?	y (F)	y
NGC 6193	CoD-48 11051		y	y	y (D2, K)	y			y (F)	y
	CoD-48 11059				y (D2, K)	?			n (F)	?
NGC 6281	HD 322676		p		p (K)				n (F)	?
	HD 153948		y	y	y (D, K)	y	y	y	y (F)	y
NGC 6383	NGC 6383 26				p	?			n (F)	?
	HD 317857		y	?	? (K)	?			y (F)	y
	NGC 6383 27				?	?			n (F)	?
NGC 6405	HD 318107		y	y	y (D, K)		y	y	y (o, F)	y
	HD 318100		y	y	y (K)		y	y	y (F)	y
	CoD -32 13119		p	y	p (K)		y		n (F)	p
NGC 6475	HD 162305		y	y	y (D, K)	n	p		n (F)	p
	HD 320764		p	y	p (D, K)		y		n (F)	p
	HD 162725	y	p		y (B, D, K)	y	y	y	n (F)	y
NGC 6633	HD 169959A		p	?	? (D, K)	p	y		y (F)	y
	HD 170054	p	y		y (B, D, K)	?	n	?	n (F)	?
IC 4725	BD-19 5044L			y	p	n	p		n (F)	p
	HD 170836		y	p	p (D, K)	?	y		y (F)	y
	HD 170860		p	p	p (D, K)	n	y		n (F)	p
Mel 227 = Cr 411	HD 190290				p	y			y (p, F)	y

Notes on individual stars (* in "member?" or "Ap?" column):

HD 19805	Composition normal, no field, not Ap star (private communication, J. Silvester)
HD 23387A	Composition normal, no field, not Ap star (private communication, J. Silvester)
HD 37022	Tycho-2 μ_δ inconsistent with Hipparcos value and with membership of this Trapezium star in Ori OB1d. Accept Hipparcos value.
CpD -20 1640	This star is Cox 40. Simbad identification and coordinates of NGC 2287 40 are incorrect.
HD 56343	Parallax and proper motions agree with cluster means, but Baumgardt et al 00 regard this star as non-member. We accept it as member.
HD 62376	Carrier et al. (1999) and Baumgardt et al. (2000) accept star as probable member. Because of 5σ discrepancy in μ_δ we consider it a non-member.
HD 65712	Star seems to be outside field considered by Dias et al 2001.
CpD -60 944B	Dias et al 2001 consider this star non-member, but both proper motions differ from cluster mean by less than 2σ. Accept as probable member.
HD 65949	Both Hipparcos and Tycho-2 proper motions have large uncertainties, but are consistent with probable membership. Disregard small membership probability assigned by Dias et al 2001.
HD 74195	Hipparcos and Tycho-2 μ_δ's inconsistent. Accept Tycho-2 value which is consistent with membership.
HD 136347	Tycho-2 μ's both completely different from Hipparcos values. Accept Hipparcos values and membership.
HD 145501	Membership assessment for ν Sco CD based on π and μ's for ν Sco AB.
HD 190290	Cluster mean motions unknown for Melotte 227.

References for "member?" column:

(B) Baumgardt et al. (2000); (Br) Brown et al. (1999); (C) Carrier et al. (1999); (D) Dias et al. (2001); (D2) Dias et al. (2002); (K) Kharchenko et al. (2005); (M) Marschall et al. (1982); (R) Robichon et al. (1999); (S) Sabogal-Martínez et al. (2001); (T) Tian et al. (1996); (Z) de Zeeuw et al. (1999).

References for "mag fld" column:

(F) Bagnulo et al. (2006), FORS1 data from Paper I; (a) Bychkov et al. (2003); (b) Brown et al. (1985; (c) Borra (1981); (d) Bohlender et al. (1987); (e) Wade et al (2006); (g) Thompson & Landstreet (1985); (h) Bohlender et al. (1993); (i) Preston et al. (1969); (j) Shorlin et al. (2002); (k) Borra et al. (1983); (l) Landstreet (1990); (m) Landstreet et al. (1979); (n) Thompson et al. (1987); (o) Mathys et al. (1997); (p) Hubrig et al. (2004); (q) Babcock (1958); (r) unpublished Musicos data, courtesy J. Silvester; (s) Kochukhov & Bagnulo (2006).

Table 3. Physical properties of stars that are probable open cluster members, probable Ap stars, and have magnetic field measurements.

Cluster	$\log t$	True DM	Star	$\log T_e$	$\log L/L_\odot$	$M/M_\odot$	Fractional age	B_{rms}
NGC 1039	8.30 ± 0.15	8.489	HD 16605	4.025	1.65	2.55 ± 0.15	0.37 ± 0.15	1700
α Per	7.93 ± 0.10	6.40	HD 21699	4.158	2.78	4.60 ± 0.20	0.70 ± 0.21	565
NGC 1662	8.64 ± 0.06	8.20	HD 30598	3.939	1.90	2.65 ± 0.15	0.91 ± 0.21	190
Ori OB1a	7.00 ± 0.10	7.63	HD 35298	4.201	2.36	4.25 ± 0.25	0.07 ± 0.02	2280
			HD 35502	4.209	2.66	4.80 ± 0.20	0.09 ± 0.03	1520
Ori OB1b	6.55 ± 0.15	8.37	HD 36485	4.290	3.27	6.80 ± 0.30	0.07 ± 0.03	3220
			HD 36526	4.212	2.54	4.65 ± 0.20	0.03 ± 0.01	2230
			HD 37140	4.190	2.61	4.60 ± 0.20	0.03 ± 0.01	580
			HD 37479	4.382	3.55	8.95 ± 0.40	0.12 ± 0.05	1910
			HD 37633	4.121	2.14	3.50 ± 0.20	0.02 ± 0.01	285
			HD 37776	4.369	3.54	8.80 ± 0.40	0.11 ± 0.04	1260
Ori OB1c	6.66 ± 0.20	8.52	HD 36540	4.196	2.77	4.80 ± 0.30	0.04 ± 0.02	275
			HD 37058	4.312	3.26	7.05 ± 0.30	0.09 ± 0.05	710
			HD 37642	4.164	2.56	4.25 ± 0.20	0.03 ± 0.02	2140
Ori OB1d	6.00 ± 0.20	8.52	HD 37022	4.595	4.95	45.00 ± 5.00	0.33 ± 0.16	205
NGC 2169	6.97 ± 0.10	10.11	NGC 2169 12	4.140	2.20	3.65 ± 0.15	0.04 ± 0.01	3410
NGC 2232	7.55 ± 0.10	7.56	HD 45583	4.104	2.04	3.30 ± 0.15	0.13 ± 0.04	2730
NGC 2244	6.40 ± 0.10	10.80	NGC 2244 334	4.204	2.35	4.30 ± 0.30	0.02 ± 0.01	9515
NGC 2281	8.63 ± 0.05	8.733	HD 49040	3.944	1.93	2.65 ± 0.15	0.89 ± 0.19	160
NGC 2287	8.32 ± 0.12	9.20	CPD-20 1640	3.908	1.16	1.85 ± 0.10	0.16 ± 0.05	155
			HD 49299	3.987	1.60	2.40 ± 0.10	0.33 ± 0.11	2720
CMa OB2	7.05 ± 0.20	8.37	HD 51088	4.097	2.23	3.41 ± 0.25	0.04 ± 0.02	90
NGC 2323	8.15 ± 0.07	9.84	HD 52965	4.090	2.64	4.05 ± 0.20	0.86 ± 0.21	65
Cr 132	7.30 ± 0.20	8.37	HD 56343	4.061	2.03	3.05 ± 0.15	0.06 ± 0.03	3610
NGC 2422	8.08 ± 0.11	8.48	HD 61045	4.114	2.47	3.85 ± 0.20	0.65 ± 0.20	430
NGC 2451A	7.70 ± 0.10	6.38	HD 63401	4.130	2.25	3.70 ± 0.20	0.24 ± 0.07	365
NGC 2516	8.15 ± 0.10	7.77	HD 65712	3.996	1.41	2.30 ± 0.10	0.20 ± 0.06	865
			CPD -60 944A	4.100	2.10	3.30 ± 0.15	0.52 ± 0.15	250
			CPD -60 944B	4.107	2.03	3.25 ± 0.20	0.50 ± 0.15	120
			CPD-60 978	4.072	1.88	3.06 ± 0.15	0.43 ± 0.13	85
			HD 65987	4.100	2.32	3.60 ± 0.20	0.64 ± 0.19	540
			HD 66295	4.045	1.65	2.60 ± 0.15	0.28 ± 0.08	490
			HD 66318	3.959	1.31	2.05 ± 0.10	0.15 ± 0.04	4385
NGC 2546	8.20 ± 0.05	9.82	NGC 2546 201	4.079	2.40	3.70 ± 0.40	0.77 ± 0.17	595
			NGC 2546 197	4.000	1.62	2.40 ± 0.30	0.25 ± 0.05	20
IC 2391	7.70 ± 0.15	5.82	HD 74169	4.009	1.43	2.35 ± 0.10	0.08 ± 0.03	130
			HD 74535	4.133	2.39	3.85 ± 0.15	0.27 ± 0.11	95
NGC 2925	8.17 ± 0.13	9.44	HD 83002	4.093	2.41	3.70 ± 0.25	0.72 ± 0.25	20
NGC 3114	8.13 ± 0.15	10.01	HD 304841	4.090	2.17	3.40 ± 0.15	0.53 ± 0.21	335
			HD 304842	4.097	2.29	3.55 ± 0.15	0.59 ± 0.23	20
vdB-H 99	7.80 ± 0.20	8.63	HD 92190	4.104	2.28	3.55 ± 0.15	0.28 ± 0.14	10
IC 2602	7.65 ± 0.20	5.91	HD 92385	4.045	1.75	2.70 ± 0.15	0.10 ± 0.05	440
			HD 92664	4.152	2.48	4.05 ± 0.20	0.27 ± 0.13	810
Cr 228	6.60 ± 0.20	12.77	Cr 228 30	4.362	3.38	8.00 ± 0.40	0.11 ± 0.05	560
NGC 3532	8.45 ± 0.10	8.47	HD 96040	4.025	1.47	2.40 ± 0.15	0.45 ± 0.13	260
Coma Ber	8.70 ± 0.10	4.70	HD 108945	3.944	1.60	2.30 ± 0.15	0.71 ± 0.21	110
NGC 5460	8.17 ± 0.10	9.34	HD 122983	4.029	1.82	2.70 ± 0.15	0.32 ± 0.10	170
			HD 123183	4.000	1.71	2.50 ± 0.10	0.26 ± 0.08	410
			HD 123225	4.090	2.38	3.65 ± 0.15	0.70 ± 0.21	20
NGC 5662	7.82 ± 0.18	10.08	HD 127924	4.121	2.58	4.05 ± 0.20	0.40 ± 0.18	10
Lower Cen Cru	7.15 ± 0.15	5.36	HD 112381	4.000	1.53	2.35 ± 0.10	0.02 ± 0.01	3400
			HD 114365	4.072	1.83	2.85 ± 0.15	0.04 ± 0.01	290
			HD 119419	4.045	1.62	2.60 ± 0.15	0.03 ± 0.01	1430
Upper Cen Lup	7.20 ± 0.10	5.73	HD 125823	4.283	3.16	6.40 ± 0.30	0.27 ± 0.08	320
			HD 126759	4.086	1.87	2.95 ± 0.15	0.04 ± 0.01	335
			HD 128775	4.079	1.76	2.85 ± 0.15	0.04 ± 0.01	345
			HD 133652	4.114	2.09	3.35 ± 0.15	0.06 ± 0.02	1120
			HD 133880	4.079	2.10	3.20 ± 0.15	0.05 ± 0.02	2300
			HD 136347	4.056	1.78	2.75 ± 0.15	0.04 ± 0.01	225
			HD 143699	4.199	2.72	4.80 ± 0.20	0.14 ± 0.04	165
Upper Sco	6.70 ± 0.10	5.81	HD 142301	4.204	2.52	4.60 ± 0.20	0.04 ± 0.01	2385
			HD 142990	4.258	2.79	5.40 ± 0.25	0.06 ± 0.02	1080
			HD 144334	4.212	2.50	4.60 ± 0.20	0.04 ± 0.01	715
			HD 145501	4.176	2.53	4.30 ± 0.20	0.04 ± 0.01	1385
			HD 146001	4.146	2.41	3.90 ± 0.20	0.03 ± 0.01	650
			HD 147010	4.111	1.92	3.15 ± 0.20	0.02 ± 0.00	4825

Table 3. continued.

Cluster	$\log t$	True DM	Star	$\log T_e$	$\log L/L_\odot$	$M/M_\odot$	Fractional age	B_{rms}
NGC 6087	7.90 ± 0.10	10.29	CPD-57 7817	4.079	2.25	3.35 ± 0.20	0.30 ± 0.09	610
			HD 146555	4.107	2.19	3.45 ± 0.20	0.33 ± 0.10	270
NGC 6178	7.15 ± 0.22	10.03	HD 149257	4.406	3.82	10.30 ± 0.60	0.61 ± 0.33	160
			HD 149277	4.348	3.71	8.75 ± 0.40	0.45 ± 0.24	2435
NGC 6193	6.69 ± 0.20	10.31	CD-48 11051	4.301	3.32	7.00 ± 0.30	0.10 ± 0.05	2400
NGC 6281	8.45 ± 0.10	8.86	HD 153948	4.025	1.86	2.70 ± 0.15	0.62 ± 0.18	195
NGC 6405	7.80 ± 0.15	8.44	HD 318107	4.072	1.92	2.95 ± 0.15	0.17 ± 0.07	4820
			HD 318100	4.025	1.64	2.50 ± 0.10	0.11 ± 0.04	345
			CoD -32 13119	3.892	1.06	1.75 ± 0.10	0.04 ± 0.02	35
NGC 6475	8.47 ± 0.13	7.24	HD 162305	4.004	1.82	2.65 ± 0.15	0.61 ± 0.22	85
			HD 162725	3.982	2.36	3.30 ± 0.20	1.08 ± 0.38	60
NGC 6633	8.75 ± 0.05	7.93	HD 169842	3.924	1.62	2.35 ± 0.10	0.85 ± 0.18	370
IC 4725	8.02 ± 0.08	10.44	BD-19 5044L	4.107	2.25	3.55 ± 0.15	0.46 ± 0.12	235
			HD 170836	4.133	2.80	4.55 ± 0.20	0.84 ± 0.22	505
			HD 170860	4.137	2.63	4.25 ± 0.20	0.71 ± 0.19	40