Bull. Astr. Soc. India (1983) **11**, 277–296

Young galactic clusters

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Received 1983 June 27

Abstract. The evolutionary nature of the young galactic clusters, from both observational as well as theoretical points of view, is briefly discussed. In addition, some of the interesting properties of the individual objects associated with these groups are mentioned.

Key words : Galactic clusters—pre-main sequence evolution

1. Introduction

The analysis of both the galactic and the globular clusters has not only paved the way for understanding the nature of the stellar evolution but also allowed us to examine critically, and interpret the nature of, some of the individual members. The overall analysis of these groups has helped in formulating :

(i) the zero-age-main-sequence (ZAMS) i.e. the initial position of the star on the H-R diagram or colour-magnitude diagram before it actually converts hydrogen into helium through nuclear reactions; (ii) verification of the initial luminosity function or the initial mass function (IMF); (iii) the overall history of heavy element production and the variation of it at the solar galactocentric distance, and also an accurate determination of (Fe/H) ratio for various clusters; (iv) the nature of the interstellar reddening law; (v) the distribution of stars along the colour-magnitude (c-m) diagram and the identification of various gaps associated with different physical processes; (vi) the calibration of the distance scale in our galaxy as well as the universe's distance scale; and (vii) stellar activity associated with individual objects.

One can divide these galactic physical groups into three broad categories on the basis of the stellar evolution theory:

(a) Old galactic clusters whose estimated age is 10^9 yr. (*e.g.* M 67, NGC 752), (b) Intermediate-age group. (Age $\approx 5 \times 10^6$ to $\approx 10^9$ yr; *e.g.* Hyades, Coma and Pleiades etc.), (c) Very young galactic clusters. (Age $\leqslant 5 \times 10^6$ yr; *e.g.* NGC 2264, NGC 6530, Orion, Taurus-Auriga complex etc.)

The analysis of the last group is the most challenging and cumbersome from both observational and theoretical points of view owing to the presence of gas, dust and other complexities associated with them. An accurate determination of the intrinsic colours of members would provide information about their positions in the c-m diagram, the nature of the reddening, blue veiling, ultraviolet peculiarities and perhaps the existence and nature of circumstellar shells.

The young galactic clusters NGC 2264 and NGC 6530 are well known for their peculiar colour-magnitude diagrams (Walker 1956, 1957; Kilambi 1977) and for the presence of Hα emission objects including many T Tauri objects (Herbig 1954, 1962). It has been difficult to relate the observed colours and magnitudes of some of their members to the effective temperatures and luminosities for several reasons.

(i) Balmer-line emission ranges from none to very strong from star to star and varies in time for a given star.

(ii) Selected metallic line emission is present to a lesser extent in some of the stars.

(iii) There is a blue veiling phenomenon which tends to fill in the absorption features increasingly toward the violet. It ranges from none to strong from star to star and is correlated with Balmer emission.

(iv) The possible presence of circumstellar dust shells around some or all of the stars further complicates the procedure for determining the V-band absorption from the observed colour excess.

2. Reddening corrections

The nature of the intrinsic colours of each cluster member and the overall properties of the group can be evaluated only when we correct the observed data for interstellar extinction and scattering. For early type stars, the estimate of interstellar extinction can be evaluated by the 'Q' technique as given by Johnson & Morgan (1953), whereas for the late type stars this can be achieved by a 'sliding fit' technique in two colour diagram (Johnson & Morgan 1953). Both these procedures assume that the behaviour of the interstellar reddening law is uniform and normal throughout the Galaxy and is well known. In addition, both these techniques can be used for only main sequence stars. Deviations from this normal nature have been encountered by Johnson & Borgman (1963) in some parts of the Milky Way. These deviations appear mainly in regions surrounded by nebulous material.

It is very essential to investigate the ratio of $E(U-B)/E(B-V)$ directly from the available data rather than assume the normal relationship. The ratio can be evaluated provided we know accurate spectral types and luminosity classifications for a reasonable sample of cluster members. The difference between the observed and the intrinsic colours will give the colour excess, and thus the ratio can be evaluated.

3. Computation of $R(= A_V/E(B-V)$

The ratio R of total to selective visual absorption can be evaluated provided we know the spectral types of a good number of objects in the group. From various cluster studies Johnson & Borgman noticed that the value of R varies from 3.1 to 7.4 and that the largest values of R tend to occur in very young clusters which have small values of $E(B-V)$. Usually, one uses the 'cluster method' procedure as given by Johnson (1965). With the assumption that all stars within a cluster are essentially at the same distance from us, the distance modulus $V-M_V$, where V is the observed apparent visual magnitude on the UBV system, should be a constant except for the effects of variable interstellar absorption within the cluster volume. To find R, we observe the *UBV* magnitudes and the spectral types of the cluster members. The spectral types give the absolute visual magnitude, M_V, and also the intrinsic colour, $(B-V)_0$ through the standard calibration. The colour excess $E(B-V)$ can be evaluated through the relation $E(B-V)=(B-V)_{obs}-(B-V)_0$. Then, one plots $V-M_V$ as a function of $E(B-V)$ and the slope of the straight line which best fits the observational data gives the value of *R*.

The above procedure is useful only if the objects of the group have already reached the main sequence. The next question arises whether the same values obtained through above procedures can be used for stars which are still in the gravitational contraction stage and have not had enough time to reach the main sequence. Certainly one cannot use the same reddening values, as these lead to inaccurate estimation of masses, luminosities, effective temperatures and evolutionary nature of these objects because of the presence of emission, blue veiling, etc.

Cohen & Kuhi (1979) have used line ratios of various elements such as H, He I, He II, Fe I, Ca I, Na I, Mg I, Cr I, G band of CH and TIO heads in order to classify T Tauri stars spectra from early to late type objects. The luminosity classification for these stars has been acheived by Cohen & Kuhi by using two useful discriminants namely λ6385 CaH absorption feature and λ5211 MgH feature. λ6385 is absent in the spectrum of M giants whereas it is highly visible in dwarfs (*M*0–*M*5). The T Tauri spectra reveal the presence of λ6385 but with a strength noticeably less than expected from dwarfs, thus indicating that these stars are of luminosity class IV. The MgH feature is useful in the range *K*5–*M*5 and a useful discriminant between classes III and V. Mould & Wallis (1977) have discussed the luminosity classification of T Tauri stars using the λ6940 CaH band head. Pesch (1972) has commented upon the high sensitivity of the λ5530 CaOH band to luminosity for stars later than M5, as it is present only in stars of class V. There is no indication of this band in many of the T Tauri stars, thus, implying a luminosity classification for some of these stars brighter than class V. The overall conclusion is that the T Tauri stars lie between the sequences from III to V class, lending some support to their empirical definition as class IV objects.

Cohen & Kuhi described a narrow band colour technique to determine A_V for individual stars. The region λ5400–6700 Å has been selected as being free of obvious excess continuum emission and the three small regions at λ5400, λ6000 and λ6700 Å show no strong emission lines. Once a star has been classified spectroscopically, the intrinsic narrow band colours are taken from the grid of spectral standards (corrected for their small reddening) and compared with the observed continuum. The value of *R* is assumed normal. Another way of investigating the value of *R* is to take these narrow band colours and to derive A_V from them using a range of values of *R* from 2.5 to 5.5. It is found that for an individual star the difference between the two values of A_V, derived separately from the (λ5400, 6700) and (λ6000, 6700) indices, is a minimum for $R \approx 3.0$ and increases monotonically as *R* increases. Therefore, as long as normal photospheres are a good representation of T Tauri stellar continua between λ5400 and λ6700 Å, the internally most consistent extinction is that for a normal value of $R = 3.10$.

The next problem is to understand where this extinction is located, that is whether it is interstellar, intracluster or circumstellar or whether it is an obscuration. If a pair of stars exhibit a common value A_V, then this is an argument in favour of extinction being intracluster or interstellar in origin rather than being circumstellar or obscuration. Conversely, in small groups that are spread over a few arc minutes, a large variation in A_V might signify either circumstellar shells or clumpiness on a scale of arc minutes in the dark cloud material. It has been noticed in the case of pairs that the value of A_V is generally the same (within the uncertainities) for separations of order less than or equal to 10″ both in Taurus-Auriga and Orion implying that the cloud material is smooth on a scale of $\approx 10^{16}$ cm in the former and $\approx 3 \times 10^{16}$ cm in the latter. For small groups containing early type stars, there are indications of higher A_V towards the O and B stars than is typical for the cooler stars. This could be evidence for either circumstellar shells around hot stars or more deeply embedded locations within the cloud for these stars of higher mass. It is very hard to decide between these two possibilities.

If we assume that these hot stars suffer extinction and possess no circumstellar infrared excess, then we can derive the value of R towards them using $R \approx 1.10 \times E(V-K)/E(B-V)$ and $R \approx 1.06\, E(V-L)/E(B-V)$. The difference between these two R values thus obtained is suggestive of contamination of extinguished starlight by, for example, thermal emission from circumstellar dust. Conversely, the agreement between these two values is an indication that these are heavily embedded stars without circumstellar dust shells. The values of R thus obtained are anomalously high indicating an intracluster origin for the obscuration. (The interstellar reddening between the sun and Taurus-Auriga or Orion is insufficient to account for the typical value of $A_V = 1$ mag derived in some cases.) Kuhi (1974) suggested that some of the reddening might be circumstellar in origin because of an apparent trend between A_V and spectral type; A_V decreasing with late spectral types. Finally, there is no reliable way of obtaining the value of A_V for the strong emission line stars which have no photospheric features.

4. Membership content

It is very essential to know whether a particular object in or around the cluster belongs to it or not. When one plots the apparent brightness against the observed colour index, some of these objects may occupy unexpected positions on the c-m diagram. The observed position of these objects can be interpreted in the following manner :

(i) The stars which lie on the right side of the H-R diagram could belong to the contracting population which is proceeding towards the main sequence. Or they could be background field stars, or objects which are in an advanced evolutionary stage. The contracting objects may or may not possess circumstellar shells around them. The existence of these shells could be confirmed through infrared and far-infrared observations. (ii) The objects which lie below the main sequence or to the left of it could once again belong to the field star contamination or be contracting objects with definite thick circumstellar shells around them. Alternatively they could be the stars which have already reached the ZAMS, but whose surrounding circumstellar shell has not been fully dissipated and still plays a major role in its evolution.

There are various methods to discern whether a particular object is a member of that group or not. The most obvious and trustworthy procedure is the proper motion analysis. If all the stars in a group are formed from a single cloud, then they would all share a common proper motion within a certain observable error limits. All such objects can be treated as belonging to that particular group. Alternatively, radial velocity measurements could be used, but this method is not as reliable as the proper motion study except in the case of normal stars.

These two procedures can be used for both main sequence objects as well as the contracting objects. But there is another procedure which can be used for membership identification, if neither proper motions nor radial velocities are available; but it is reliable only for the main sequence objects. Both the contracting objects as well as the stars which are slightly evolved from the main sequence cannot be identified as members definitely with this procedure. This technique is usually known as 'Walker's technique' (1965). In this, the probability of membership for stars is judged by using the V_0, M_V (ZAMS) diagram. The derivation of the unreddened apparent visual magnitudes and the M_V (ZAMS) can only be made after applying interstellar reddening corrections. The absolute visual magnitudes can be determined from the ZAMS calibration given by Johnson & Iriarte (1958) and FitzGerald (1970) using its intrinsic $(B - V)$. Stars within the domain defined by the boundary lines are considered probable members. These limits were set by assuming that typically cluster members will not lie more than $0^m.5$ above the line $V_0 = M_V +$ const and that duplicity will not brighten stars more than $0^m.75$. These limits were set by Walker arbitrarily from the studies of well known clusters, which have very reliable proper motion data.

5. H-R Diagrams

The effective temperatures and bolometric luminosities obtained from observed spectral types and energy distributions can be used to construct H-R diagrams for several associations and young cluster groups. The diagrams thus obtained are almost identical for Taurus-Auriga complex, the Orion association, NGC 2264, NGC 7000/IC 5070 and Rho Ophuchi. The diagrams show that all these groups have a well developed main sequence and, in addition, majority of the low mass stars lie above the main sequence and are still in the gravitational contraction stage. The stars which lie above the main sequence and to the right side of the H-R diagram can be divided into three broad groups on the basis of their observed properties. They are (i) T Tauri objects which are less massive contracting objects, (ii) Herbig's Ae and Be type objects which are more massive than T Tauri objects, and (iii) strong emission objects.

(i) *T Tauri objects*

T Tauri objects are characterized by the following properties :

(a) Balmer lines show emission in varying degrees.

(b) Observed emission in the H and K lines of Ca II as well as in Fe I, Fe II, Ti II and [S II] is seen in varying degrees.

(c) [O I] at λ6300 and λ6363 are in emission.

(d) The absorption line of neutral lithium at λ6708 is rather strong.

(e) Underlying spectral types indicate that between late F and middle M, the photospheric lines are often broad. The broadening of the lines has usually been interpreted as arising from rotation (Herbig 1957) and macroturbulence.

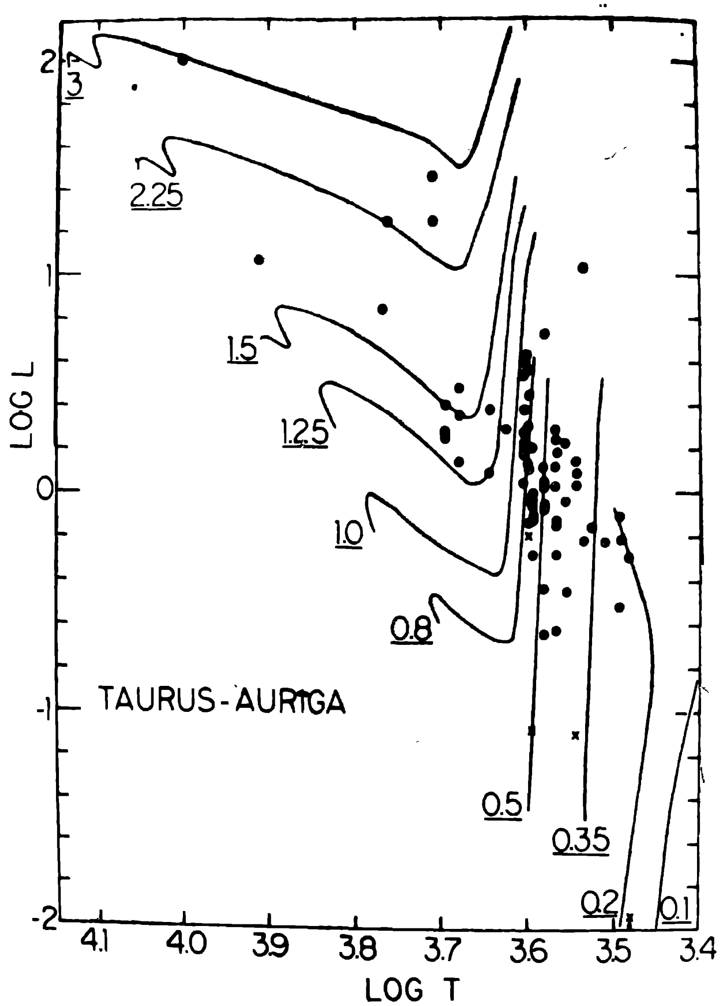

Figure. 1. Hertzsprung-Russell diagram for pre-main-sequence stars in the Taurus-Auriga region. Observed stars are represented as dots, the position of each being determined in the vertical direction by its luminosity and in the horizontal direction by its surface temperature. The curves are evolutionary tracks—paths that a star follows as its luminosity and temperature change with time, according to the equilibrium theory. (Cohen & Kuhi 1979).

(f) P Cygni type profiles have been observed in the Balmer lines and in the H and K lines of Ca II. YY Orionis, a subgroup of T Tauri objects studied by Walker (1972), is an exception.

(g) A majority of these objects exhibit an ultraviolet excess and in some cases a blue continuum also. The UV excess appears to be best interpreted as Balmer continuum radiation. However, at wavelengths shortward of 4200 Å, many T Tauri stars are observed to exhibit a steep rise in their spectral energy distribution. This is often accompanied by the apparent weakening or complete disappearance of photospheric lines. This washed out appearance is described as *veiling*. In some cases, no photospheric absorption features are visible even in the red indicating that the veiling phenomenon extends to atleast λ6500 Å. No entirely satisfactory explanation of this effect has been proposed.

(h) Many of these objects exhibit optical polarization in varying degrees. Hunger & Kron (1957) first reported optical polarization for T Tauri itself. Since then, little work was done on this question until Serkowski (1969a, b) and Breger (1974) surveyed a few objects of this class. The polarization appears to be variable with time indicating an intrinsic origin and in some cases reaching as high as 12% (S. E. Strom, personal communication).

(i) Many of these objects exhibit infrared excesses. Mendoza (1966, 1968) was the first to report this unusual characterstic of the T Tauri stars. Some, though not the majority, of these stars appear to show evidence of a $10\,\mu$ emission bump (Strom *et al.* 1972) and a sharp rise in energy distribution around $20\,\mu$ (Cohen 1973a, b).

(j) Observed irregular variability at optical wavelengths. Some of these stars have amplitudes as large as five magnitudes, but more typically of the order of one or two magnitudes are observed. No evidence of periodicity has been found (Plagemann 1970). In one case (V 1057 Cygni), a six magnitude rise in a period of less than 100d has been observed. This rapid rise in luminosity is similar to that observed in FU Orionis (Herbig 1966), a star presumably analogus to V 1057 Cygni.

(ii) *Herbig's Ae/Be type objects*

These are early type objects analogues to the T Tauri stars but more massive and hence pre-main-sequence (PMS) objects (Herbig 1960). The criteria for the identification of this class are:

(a) emission lines in the spectrum; (b) association with regions of obscuration by dust clouds; (c) illumination of reflection nebulosity; (d) irregular variability; and (e) infrared excesses (Allen 1973). (f) Loren *et al.* (1973) have detected CO emission at the locations of these objects.

(iii) *Strong emission objects*

These are very cool objects analogues to the T Tauri stars but are in much earlier stage of pre-main sequence evolution. They exhibit very strong Hα emission and usually lie on convective tracks in the H-R diagram. Their observed distribution does not correspond to any published dynamical evolutionary tracks. Most Hα emission stars are of late K spectral type and have luminosity class between III and V. The latest spectral type observed so far is M 5.5 and the estimated age of the youngest object is 10^4–10^5 yr. Radii and masses derived from the H-R diagram lie between 1 and 5 $R_\odot$ and 0.2 and 3 $M_\odot$ respectively. The observed correlations between Hα, Fe II and [O I] emission with He I suggest a chromospheric origin or atleast a common mechanism of origin for the emission spectrum. General correlations of the same emission lines with bolometric luminosity suggest a decrease in emission activity with increasing age. It is observed that the emission line activity dies down with decreasing luminosity (and hence, with increasing age) and this decay is shown more strikingly by [O I], Fe II and He I than by Hα. This in turn suggests that Hα emission may have contributions from more than one source.

A possible explanation for the general decrease of emission activity with decreasing luminosity may be found from the evolutionary tracks. The stars on the Hayashi tracks are fully convective at high luminosity and become progressively less so as they evolve, *i.e.* the fractional depth of the convective zone decreases with time. If the convective zone were the primary source of the mechanical energy that powers the chromosphere and the stellar wind, and hence the excitation of the emission lines, then the observed correlations would be more or less expected. However, the connection between bolometric luminosity and collisionally excited [O I] emission through such a stellar wind is a very tenuous one, although plausible; and no calculations have yet been made to check its feasibility. Dynamical models such as those of Appenzeller & Tscharnuter (1975) have a very different structure along the vertical track and do not show the systematic decreases in convective energy transport of the Hayashi track. In this case, the energy source for the excitation of the emission spectrum could be the kinetic energy of the infalling material, which is also expected to decrease with increasing age.

The correlation between Hα emission and the excess over photospheric flux in the near-infrared (K-L) index shows a weak trend indicating that both Hα emission and thermal emission by hot grains diminish with age, or that some of the hot dust forms within the stellar winds as indicated by Hα emission. Thus, in general the emission line fluxes for the strong line stars confirm the trends shown by the less extreme stars.

The estimation of masses and ages for individual stars can be obtained after adopting specific evolutionary calculations for contracting objects that yield theoretical tracks in the H-R diagram. The noticeable clumping of stars for Taurus-Auriga complex provides a test for the selection of tracks to be used (*e.g.* dynamical versus convective-radiative tracks). The clump (between log T_{eff} of 3.5 and 3.6 and log L of $+0.6$ to -0.4) cannot be matched by Larson's (1972) dynamical calculation at ages when the stars are expected to be visible, *i.e.*, his theoretical stars do not occupy the observed region of the clumping as visible objects. The only exception noticeable is that of Larson's 0.25 $M_\odot$ theoretical model, which becomes visible in the low luminosity part of this clump. On the other hand, the clump seem to lie along the Hayashi convective tracks. The recent calculation for 1 $M_\odot$ by Appenzeller & Tscharnuter (1975) with a better treatment of the radiation flow through the accretion shock produces a lower luminosity object (compared with that of Larson's) which converges to the convective track. The timescales are similar but the star appears redder by 10–20%. However, the changes in location of a model star during this phase of evolution are much larger than this due to uncertainities in the initial conditions (Larson 1969; Narita *et al.* 1970). The convective radiative tracks used here are adopted from various models which are consistent in their treatment of convection (Bodenheimer 1965; Iben 1965; Grossman & Graborke 1971; Taam 1977). The stars which are slightly in an advanced evolutionary stage continue their evolution towards the main sequence along the radiative tracks. The infrared and far-infrared photometric observations showed that some of these objects possess circumstellar shells around them The stars which lie below the main sequence and to the left of it can be interpreted as contracting objects with very thick circumstellar shells containing both dust and gas. The outflowing radiation is being absorbed by the dust component of the shell and reradiates in the far-infrared wavelength regions, making the object much fainter in the visual region. At the same time, the contribution from the gas emission makes the object much more hotter, and thus the net effect of both dust and gas is to place the star to the left or below the main sequence.

6. Masses and ages of contracting objects

Masses and ages for individual stars can be obtained from their location in the H-R diagram with respect to the adopted tracks and isochrones. In addition, the masses for contracting stars can be derived from the fundamental relation

$$\text{Log } M/M_\odot = \text{Log } g/g_\odot + \log L/L_\odot - 4 \log T_e/T_{e\odot},$$

where M, T_e, g and L refer to the mass, effective temperature, surface gravity and luminosity of the star and $M_\odot$, $T_{e\odot}$, $g_\odot$ and $L_\odot$ are the corresponding quantities for the sun. Thus, in order to determine the mass of an object, the value of L, T_e and g must be obtained. McNamara (1975) has determined masses for nine pre-main sequence objects in NGC 2264 by employing flux measurements extending from 0.36 μ to 3.4 μ in conjunction with the model atmospheres. The resultant masses are in good agreement with the theoretical predictions although there appears to be a systematic trend in that the stars located near the main sequence are more massive than expected from evolutionary calculations of nonrotating stars (Iben 1965). A possible explanation for this behaviour can be seen in the calculations by Bodenheimer & Ostriker (1970) on the evolution of differentially rotating stars: When a differentially rotating star is near or on the main sequence, the mass predicted on assuming zero rotation is too low. The predicted mass difference between rotating and nonrotating models increases as the angular momentum of the star is increased at a given orientation and location in the H-R diagram. However, other parameters such as inclination, degree of differential rotation and geometrical distortions produced by rotation make a detailed comparison difficult.

As mentioned earlier, the age of a contracting object or the overall age of the group can be inferred from theoretical isochrones. Apparently, all the stars in a cluster do not have exactly the same age. The estimated age spread in NGC 2264 is almost 10 Myr (Warner *et al.* 1977; McNamara 1975; whereas the age spread in NGC 6530 is about 3 to 7 Myr (Kilambi 1977). This fact is important in the study of the properties of very young clusters. In particular, it is clear that the IMF of a very young cluster varies with time, if the age spread depends on stellar mass. Iben & Talbot (1966) and Williams & Cremin (1969) have raised the question "Do low mass stars form first ?". No conclusive answer has yet been available for this question.

The distribution of stars in major associations according to spectral type and age are suggestive not of coeval (on a timescale less than from $\approx 10^4$ to $\approx 10^6$ yr) star formation but rather of an initial jump in activity followed by a continuous production of stars. The age histogram for Taurus-Auriga and Orion shows that the number of stars that are closely nebulous (*i.e.* intimately associated with local nebulosity as opposed to being projected upon general background nebulosity) decreases as the age of the sample increases, especially in Taurus-Auriga region. One might conclude that these closely nebulous objects are preferentially the youngest. Another interesting noticeable observation is the proportion of stars having continuous spectra and strong lines among the various stellar groups. Indirectly, this proportion is the ratio of the number of stars on convective tracks to that on radiative tracks. From this study, one can infer that Taurus-Auriga can be described as the most youthful and vigorous of the various complexes and NGC 2264 as the oldest and the least active.

7. Stellar activity during pre-main sequence evolution

In recent years, considerable effort has been expended in understanding the role of stellar activity on the main sequence and the subsequent evolutionary stages. The most obvious indicators of stellar activity are the Ca II K chromospheric emission, variability, stellar rotation, transition region emission lines in the far ultraviolet and coronal x-ray emission. Skumanich (1972), using the work of Wilson and Kraft, showed that the flux in the Ca II K-line, the lithium resonance line strength and the stellar rotation rate all decay as (age)$^{-1/2}$. The relationship seemed to hold good from stars as young as the Pleiades to stars as old as the sun. Can these conclusions be extended to the pre-main sequence (PMS) objects which are considerably younger ? How far back in time can we extrapolate the Skumanich relation ? Some answers to these questions can be found from a study of young galactic clusters.

A proper survey of the Ca II K emission strengths in young groups has not yet been done. But low resolution scanner data obtained by Kuhi (1973) do indicate that in NGC 2264, the non-Hα-emission stars obey the Skumanich relation. However, it is noticed that the Hα emission objects (T Tauri stars) have Ca II K emission fluxes which are too strong for the (age)$^{-1/2}$ formula if one assumes that all the stars are of the same age. But, if the spread in the age implied by the location of stars on the H-R diagram were taken into account, some of these youngest stars would have Ca II K emission in agreement with the Skumanich relation, but the range is almost a factor of ten. Duncan (1981) also reached similar conclusion from the study of single T Tauri objects. Thus a clear cut conclusion is not yet possible; a large sample should be studied.

A small number of T Tauri objects examined with International Ultraviolet Explorer (IUE) indicate the presence of Mg II, Si IV, C IV and He II lines from the early analysis by Gahn *et al.* (1979). The strength of these lines seems to vary from star to star. At the same time, the strength of some of these lines is much stronger than the others; and sometimes the line is fully absent in a particular star. This kind of behaviour in line emission implies either a missing or a very weak transition region in which the temperature is less than 2×10^5 K or the presence of a chromosphere. Boesgaard (1981) has made an attempt to put the observed line fluxes directly on the Skumanich relation with some degree of success.

Herbig (1957) found $V \sin i$ ranging from 20–65 km s^{-1} for four bright T Tauri stars. Walker (1956) had noticed the broad appearance of the absorption lines in some of the cooler stars of NGC 2264 which would correspond to $V \sin i \approx 100$ km s^{-1} at that dispersion. Vogel & Kuhi (1981) studied a large number of T Tauri stars in Taurus-Auriga complex and in NGC 2264 including some non-Hα emission objects. Their results indicate that most low mass stars ($M \leqslant 1.5\ M_\odot$) did not have measurable velocities even though the minimum observable limit for $V \sin i$ was 25 km s^{-1}. Even among the Hα emission stars, it had been noticed that a measurable velocity was an exception rather than the rule. The stars measured by Herbig were at least one $M_\odot$ in mass and hence more massive than the stars at the break of the main sequence rotation curve. Hence, Vogel & Kuhi concluded that

T Tauri stars are not rapid rotators ($\leqslant 25$ km s^{-1}) except a few stars of mass > 1.5 $M_\odot$. Nor are non-Hα emission stars rapid rotators unless their masses > 1.5 $M_\odot$. In other words, they have suggested that the break in rotational velocity is already present in the PMS stars as soon as the stars produce a photospheric absorption spectrum. The extrapolated relation of Skumanich should give a velocity of $\approx$ 80–125 km s^{-1} for T Tauri stars. Vogel & Kuhi were not able to determine $V \sin i$ for the most extreme emission line stars which have no photospheric absorption lines. It would be most desirable to obtain data with a high velocity resolution, eg. ≈ 5 km s^{-1} to test the modified Skumanich relation and to study the changes in rotational velocity that occur along the evolutionary tracks.

Ku & Chanan (1979) and Chanan *et al.* (1979) detected 23 discrete sources in Orion with the *Einstein* imaging proportional counter. Twenty five nebular variables fell within the error circles of the x-ray emission regions and the observed x-ray fluxes were in the range $10^{30.6}$ to $10^{30.9}$ ergs s^{-1}. The stars identified could be classified as $\approx 50\%$ T Tauri type and $\approx 50\%$ flare stars or $\approx 75\%$ irregular variables and $\approx 25\%$ flare stars depending on the exact criteria used. Followup work by Chanan shows that L_X is as large as $10^{31.5}$ ergs s^{-1} for F and G stars in Orion, *i.e.* ≈ 30 times the maximum luminosity observed in the Hyades. Also, the fraction of stars detected above the sensitivity limit ($L_X \approx 10^{30}$ ergs s^{-1}) in Orion is 25 to 50%, which indicates a much higher median L_X than in the Hyades. The obvious conclusion is that the x-ray luminosity decreases with increasing age, in quantitative agreement with the behaviour noted for the main sequence objects. X-ray emission may be due to (i) presence of a hot corona as in main sequence stars, (ii) shock heating of the interstellar matter by mass outflow, or (iii) shock heating of the upper stellar atmosphere by infalling material. They have also suggested that nondetection of x-ray emission is due to variability, or the emitted x-ray flux being below the threshold sensitivity, or due to nonspherical emission, *e.g.* beaming as predicted by accretion models. More observational data are required for a better understanding of the nature and origin of the x-ray emission.

8. Abundance variations in young stellar groups

One of the most accurate ways of determining metal abundances for larger samples of stars is probably through observations of the Strongren m_1-index. Crawford (1975) has defined the metal abundance indicator δm_1 as the standard Hyades m_1-value corresponding to the β-index of the star *minus* the observed m_1 value of the star. The reddening-corrected metal abundance indicator is designated as δm_0. Nissen (1979) showed that m_1 is an excellent indicator of metal abundances for F-type stars and applied the procedure to a number of clusters. His analysis showed that the average metal abundance differences between various clusters are fairly small, but the differences are significant to a high degree of confidence due to the very small errors associated with the average values of Me/H (*i.e.* logarithmic metal-to-hydrogen ratio of a star minus that for the sun). He concluded that the metal-to-hydrogen ratio for Hyades and Praesepe is about 40% higher than for Coma and NGC 752, whereas the ratio for α Persei and Pleiades lies somewhere in between. He also concluded that the metal abundance differences between various groups are not related to the ages, but could be related to the large scale metal abundance gradient in our Galaxy found by Mayor (1976) and Janes (1979).

Spectroscopic determinations of the metal abundances of stars in clusters other than the Hyades are rare due to the faintness of the stars. Nearly all the studies have been based on the observations of the equivalent widths of rather strong lines, which make the results less reliable, because the strength of these lines may be affected by differences in microturbulence and/or the damping parameter. Barry *et al.* (1979) have analysed the spectra of a group of NGC 2264 members selected on the basis of their normal and near-normal appearance. The measured equivalent width of various Balmer lines and metallic features are directly compared with those of the Hyades and Coma cluster members. It has been noticed that for a given Balmer line strength, the emission-free NGC 2264 members have metallic line strengths which are weaker than those of the Hyades members and almost equal to those of the Coma members. It is concluded that the metallicity of NGC 2264 is the same as that of the Coma cluster within a few hundredths dex. The data presented in table 1 show no evidence of an increase in the heavy element abundance at the sun's galactocentric distance during the last 5×10^9 yr.

Table 1. Fe/H ratio and the approximate age for various objects

Object	Fe/H	Approx. age (yr)
Sun	0.00	5×10^9
Hyades	0.01	10^8
Coma	—0.15	10^8
M67	—0.08	5×10^9
NGC 2264	—0.15	$0 - 10^7$

The difference in metallicity between the members of the moderately young Hyades and Coma clusters exceeds the difference between the members of the oldest (M 67) and the youngest (NGC 2264) clusters. Whereas there is ample evidence in the literature that metallicity depends strongly on location in the Galaxy, the data for open clusters do not support metallic enrichment in the galactic disc within 1.5 kpc of the sun.

Spectroscopic determinations of the helium abundances of B-type stars in young clusters and associations are relatively few. The view that the helium abundance has a universal value of $Y = 0.25$, close to what is predicted by the standard big bang theory, seems to be generally accepted. But there is growing evidence that the variation of helium abundance of the order of $Y \approx 0.1$ occurs even within our Galaxy. Peterson & Shipman (1973) investigated helium abundances of B stars in clusters and associations and found that the two associations Sco OB II and Lac OB I contain a normal helium to hydrogen ratio of ϵ (He) $\approx$ 0.10 (by number), whereas the cluster NGC 2264 situated 800 pc away from the sun in approximately anticentre direction has ϵ (He) $\approx$ 0.08. The difference was estimated to be significant at the 3 σ level. Nissen (1974, 1976) studied extensively helium abundances of main sequence B stars in young clusters and associations through a narrow-band index, I (4026), of the He I λ4026 line. A model atmosphere analysis showed that for main sequence stars in the spectral range B0–B2, the relation between I (4026) and β is insensitive to the differences in gravity but sensitive to differences in the helium-to-hydrogen ratio. The analysis showed that stars in Sco-Cen, NGC 6231, Lac OBI fall along the relation between I (4026) and β defined by nearby field stars, indicating that all these groups have the same helium abundance. On the other hand, the stars in h and χ Persei and Cep OB III were found to deviate systematically from the field star relation indicating that their helium-to-hydrogen ratio is a factor of 1.7 lower than for the other clusters. The absolute values of the helium-to-hydrogen ratio is not well determined, but if we normalize the average helium abundance of nearby field stars to be ϵ (He) ≈ 0.10, corresponding to $Y \approx 0.28$, then Cep OB III and h and χ Persei are found to have a helium to hydrogen ratio of ϵ (He) $\approx 0.060 \pm 0.004$, corresponding to $Y \approx 0.19 \pm 0.01$. Thus the helium difference is close to 0.1 in Y.

The analysis also showed that helium differences are not related to the ages of the clusters and the deficiency in h and χ Persei may be related to the radial helium abundance gradient in our galaxy found by Churchwell *et al.* (1978) from a study of helium and hydrogen radio recombination lines. However, the deficiency found by Nissen is larger than predicted from the gradient. Furthermore, the galactocentric distance of the other helium deficient group, Cep OB III, is the same as the distance of Lac OB I and Ori OB I, which were found to have a normal helium abundance. This suggests that in addition to a possible galactic helium abundance gradient, local variations of up to 0.1 in Y may occur. Nissen (1976) found the variations in the helium-to-hydrogen ratio among the member stars to be less than $\approx 20\%$ for a given cluster or association.

9. The nature of the mass spectrum or IMF

The initial mass function (IMF) of stars is one of the fundamental parameters for the study of galactic evolution. The nature of the IMF is not well known in the distant parts of the Galaxy due to lack of observations; and even in the solar neighbourhood, the IMF is rather uncertain for all mass ranges (Audouze & Tinsley 1976). According to Hoyle (1953), Rees (1976) and Silk (1977), the IMF could result from an opacity-limited hierarchical fragmentation process. In a collapsing cloud, fragments may themselves lead to further breakup, if the gas can lose enough internal energy for the Jeans mass to decrease as the density rises. In addition, if cooling were efficient, collapsing clouds would further undergo hierarchical fragmentation into progressively small masses. This process would terminate only when individual fragments become opaque enough to trap their radiation. Hence, the physical situation is not so simple and the actual models of cloud fragmentation should take into account the accretion mechanisms (Larson 1978) and also the inelastic collisions between fragments (Silk & Takahasi 1979).

Salpeter (1955) used the observed luminosity function for the solar neighbourhood and theoretical evolutionary lines to derive an IMF which may be approximated by a power law :

$$\xi\,(\log M) \approx M^{-x}$$

or

$$n(M) \approx M^{-(x+1)},$$

where ξ (log M) is the number of stars per unit logarithmic mass interval and $n(M)$ the number of stars per unit mass interval. The value of the exponent x is 1.35 for masses between 0.4 $M_\odot$ to 10 $M_\odot$. Audouze & Tinsley give

$$x = 2.6 \quad \text{for } 2 \leqslant M/M_\odot \leqslant 20,$$

$$x = 0.25 \quad \text{for } 0.4 \leqslant M/M_\odot \leqslant 1,$$

$$x = 1.3 \quad \text{for } 0.2 \leqslant M/M_\odot \leqslant 0.4.$$

Mayor & Martinet (1977) give

$$x < 1 \quad \text{for } 1 \leqslant M/M_\odot \leqslant 2.$$

In connection with the formation of open clusters, it is important to know whether the observed mass spectrum of open clusters has the same slope as the IMF for the solar neighbourhood. From the studies of Jaschek & Jaschek (1957) and Sandage (1957), one may conclude that there is a satisfactory agreement between Salpeter's IMF and the IMF of clusters, at least in the range $1 \leqslant M/M_\odot \leqslant 10$. Variations of x value have been noticed in open clusters especially for stars $M/M_\odot \leqslant 1$. But the position of this turnover may not be considered as definitely established because the determination of the luminosity function of clusters is subject to many observational biases. According to Kholopov (1969), every cluster exhibits at least two main regions, a nucleus and a corona and the ratio of corona radius to that of nucleus radius ranges from 2.5 to 10. Generally, only the cluster nucleus is taken into account for the cluster studies and this biases the cluster luminosity function because in almost all cases the cluster coronae contain a larger percentage of low-mass stars than the clustes nucleus. Moreover, the study of the low-mass IMF in very young clusters is difficult because many of these stars are in the pre-main sequence evolutionary stage for which mass attribution is very difficult.

The best way to investigate this mass spectrum for young aggregates is to consider the cumulative mass function, $N(m)$, the number of stars of mass greater than 'm'. The value of x can be obtained by using

$$d(\log N)/d(\log m) = -x.$$

The most reliable values of x are as follows :

Taurus-Auriga	x ranges from	1.2 to 2.5	with best central fit at	1.5
NGC 2264		0.8 to 2.0		1.8
Orion (all stars)		1.3 to 1.5		1.35

10. Stellar duplicity

The determination of the luminosity function or the mass spectrum is not complete till the components of double or multiple stars are included. J. C. Mermilliod (1979, personal communication) reconstructed the luminosity function of the Preasepe by assuming that the vertical dispersion above the main sequence in the observational H-R diagram is entirely due to duplicity and noticed that the mass spectrum is much steeper than before. Thus, the behaviour of the IMF in the Galaxy depends also on the rate of binary stars produced in that region. Our knowledge of this rate is extremely incomplete particularly due to large scale variations of low mass binaries in the Galaxy. Crampton *et al.* (1976) found that the percentage of known or suspected binary stars in six open clusters (α Per, Pleiades, M39, IC 4665, NGC 2516, and NGC 6475) is approximately constant (40–50%). In any case, if the IMF is used to study the processes of cloud fragmentation, the frequency of secondary components in multiple systems ought to be taken into account.

One expects to find a significant number of binary systems among the T Tauri stars, if these are indeed the predecessors of solar type stars (Abt & Levy 1976). Repeated spectroscopy of a few bright T Tauri stars, whose lines display systematically variable radial velocities, did not provide any information about the binary nature (Herbig 1977). It would be difficult to detect the duplicity of a system with mass ratio close to one, because of its spectroscopic compositeness One needs a system with very different component masses to know the nature of the secondary. Cohen & Kuhi (1979) in their survey of T Tauri stars in very young groups have detected a number of close companions with Lick TV system. If these are true visual binaries, then one would expect the two to have the same age as well as extinction. The survey indicates that 53% of the close pairs in Taurus-Auriga and 42% in Orion satisfy the comparable age criterion. Even though the existence of such binaries if confirmed by further observations would affect the completeness of the sample of the lowest mass, it does not significantly perturb plotted points on the H-R diagram. Only binaries with mass ratio close to unity would affect the H-R diagram and no information on these is yet available. In order to overcome the contamination of the optical pairs, Cohen & Kuhi suggest a common value of A_V alongwith a separation of 10″ as the best criteria for the identification of binaries in Taurus.

11. The degree of star formation

A figure of great interest to the theorists is the efficiency of the process of star formation, *i.e.* the fraction of the total gas in a region that is converted into stars. The crudest estimate one can make is to add together the masses of all the young stars that have been observed and then divide with the total hydrogen mass in the entire complex. The value thus obtained is a lower limit but still presents a general picture. In general, the star formation seems to be about 10% efficient. For smaller condensations containing only low-mass objects, the efficiency is $\approx 5\%$. Given the growth rate and the efficiency of star formation in an aggregate, we can estimate the longevity of dark clouds to be approximately 10^7 yr assuming that the present growth rates are maintained until the exhaustion of the gas. The space densities of observable young stars in these various regions are 1 to 30 stars per cubic parsec, based on the above considerations. The latter is one or two orders of magnitude greater than the density of field stars in the solar neighbourhood.

12. Pre-main sequence evolution—general theoretical picture

It is generally accepted that a star forms when an interstellar cloud of gas and dust collapses under its own gravity. Initially, the cloud falls freely toward its own centre. The matter at the centre collects into a core in which the collapse is gradually replaced by a slow contraction, because the pressure in the gas nearly balances the force of gravity. The outer part of the cloud continues to fall in but remains opaque, so that the embryonic star is hidden in a cocoon until the process is nearly complete. In the core, contraction continues until the temperature rises to the point where nuclear fusion can occur, at which time the young star is said to have reached the main sequence.

Within this broad framework, however, there is no generally accepted theory of the details of the star formation. Several model calculations disagree significantly on the size and the luminosity of the protostar at any given time. Particularly important is the end of the major accretion stage, when the infant star emerges from its cocoon and becomes observable in visual light. Various theories predict values for its radius at this phase ranging from 2 to 5 times the solar radius for a star of one solar mass. This disagreement prevents an interpretation of the observations of very young objects such as T Tauri stars. The existence of a valid theory of star formation, on the other hand, would allow these stars to be identified with an evolutionary stage and their masses can be estimated. Pioneering work in the theory of star formation was done by Chushiro Hayashi and his co-workers and Louis Henyey and his co-workers (Hayashi 1961; Henyey *et al.* 1959). Both these groups studied mathematical models of the slow contraction (under different conditions) of nearly equilibrium gas spheres at constant mass. These models took into account each stage in a contracting star and the detailed physical properties at each point within it.

The path that a protostar follows on the temperature-luminosity diagram is called the Hayashi track. Protostars are generally thought to be objects with surface temperatures less than 3000 K which are evolving through the region of the H-R diagram that was found by Hayashi (1961) to be forbidden for stellar models in hydrostatic equilibrium. Two important mechanisms cause hydrodynamic instability for cool objects in this region. The first mechanism occurs at a stage when temperatures are about 10 K and densities are 10^{-13} g cm^{-3} or less. At this point, the protostar is transparent to its infrared radiation so that much of the gravitational energy released is lost through radiation and does not result in internal heating. The other mechanism is the dissociation and ionization of hydrogen and consequent reduction of the adiabatic exponent $\Gamma_1 \equiv (d(\log P)/d(\log \rho))_s$ to values less than 4/3, in the temperature range 1800 K to about 3×10^4 K. Collapse occurs because the buildup of internal pressure is not as rapid as that of gravitational forces and proceeds until enough gravitational energy has been released to dissociate and ionize the hydrogen and heat the object up to central temperatures of the order of 10^5 K, after which hydrostatic equilibrium is possible. The high temperature results from the conversion of potential energy into energy of motion for the atoms, of which temperature is a measure. Because of the great difference in the temperature between the central parts of the protostar and its outer ones, the interior of the gas cloud is unstable and all detailed calculations show that convection must occur. In this regard the protostar is like a kettle of boiling water. For a contracting protostar, the very hot central temperature leads to convective boiling so that the entire star consists of zones that are circulating outward from the centre, bringing the cooler portions into the inner regions where they are heated and expelled again by the convective motion. The convection in a protostar is a characteristic of the initial phases on the Hayashi track of a star. Because of it, the temperature of the star at the surface remains fairly constant for a considerable amount of time. In the first few million years, the temperature changes only by 1000 to 2000, thus changing surface temparature from 3500 K to roughly 4500 K in a typical case.

During this period, the luminosity decreases because of the continuing contraction of the star. Because the total luminosity of a star depends upon the amount of the star's area that is radiating away energy and because the surface area of a star depends upon its size, the total luminosity must decrease unless of course, internal sources of energy can heat up the star. During these initial phases of nearly even temperature, the luminosity decreases by a factor of about 1000. For a protostar with a mass about that of the sun, the brightness changes by about 500 times the solar luminosity in about 2 or 3 years. After about 10 million years of contraction, this brightness decreases to about half the solar luminosity. After thousands of years (for a heavy star) or millions of years (for a light star), the configuration of mass within the protostar becomes such that convection gradually stops in the centre. The transportation of the energy from the hot centre to the outer parts can no longer occur through the circulation of the hot gas outward from the centre. Instead the energy moves outward from the centre by means of radiation. The star has reached sufficiently high densities by then so that a photon of light has a difficult journey outward through the object. The light from the hot central areas is transmitted outward slowly, because it must interact with each different layer of atoms. At a particular time in the contraction phase, this radiative stage sets in rapidly and the result is a stabiliziation of the luminosity of the star. From then on, its luminosity stays nearly the same, but the temperature of the entire star increases owing to further contraction. Thus, an object which for the first million years of its life appears as a dull red globe gradually becomes bluer as it nears the main sequence. Because of the outward-increasing pressure, which is the result of the high temperature in the centre of the star, the rate of contraction is slowed down in the later phases on the Hayashi tracks than in the beginning. In only 1 or 2 years, the gas cloud contracts from an immense object to the size of a planetary orbit. From then on, the contraction becomes slower so that the entire convective phase of contraction occupies millions of years. For a star with the mass of the sun, convection ends after an interval of 10 Myr and the radiative phase of contraction onto the main sequence takes an additional 17 Myr. The rate of evolution is much faster for heavier stars than for lighter ones.

Larson was the first to treat the complete problem of the star formation including rapid fall. Since spherical collapse is the most idealized case, calculations with spherical symmetry have been carried to a stage where a 'stellar' object is formed. The main general features of a spherical collapse are: (i) a runaway increase in central condensation, (ii) formation of a stellar core, and (iii) infall of the remaining envelope into the core through accretion shock. Larson's (1969, 1973a, b) calculations have shown that the radius and luminosity of the visible stellar core are sensitive to the assumed initial conditions, especially, the initial density. If the initial density for 1 $M_\odot$ is such that the protostar is marginally Jeans unstable, the visible pre-main sequence stellar core is relatively small ($\approx 2\ R_\odot$) and is not very luminous ($L \approx$ 10–20 $L_\odot$). However, if the initial density is significantly larger than that required for the mass to be Jeans unstable, the visible stellar core is much larger ($\approx 50\ R_\odot$) and more luminous ($\approx 10^3\ L_\odot$). The reasons behind the large difference in radius and luminosity have been explained in detail by Woodward (1978).

Briefly, if the collapse is adiabatic and there is no appreciable loss by radiation, the binding energy of the final state is equal to the energy required to dissociate H_2

and ionize all the hydrogen and helium. The estimated final radius based on this energy argument is about 50 $R_\odot$ for 1 $M_\odot$, close to the value obtained by Narita *et al.* (1970). In the calculation of Larson, the free fall time is so long that considerable energy is being lost by radiation, requiring more release of gravitational energy and hence a smaller radius in the final state. An additional consequence of the assumed initial density is the dependence of the PMS evolution on mass. If the assumed initial density is near the Jeans density, the initial equilibrium state of the PMS object depends strongly on its mass, whereas if the assumed initial density is large enough to ensure that collapse is adiabatic, the initial equilibrium state always lies at the top of the Hayashi track for the corresponding mass. A critical aspect of the evolution from protostar to a PMS object involves shocks formed by accretion of material onto an equilibrium stellar core.

The calculations showed that for an object of 2 $M_\odot$, the evolution is well into the PMS radiative contraction phase at the equilibrium point. The equilibrium core reaches the main sequence and starts nuclear burning while it still has an optically thick infalling envelope for masses $\geqslant 5\ M_\odot$. If the collapse starts at densities above the Jeans limit, the mass above which this effect occurs increases because of the shorter free-fall time. At masses below 1 $M_\odot$, the free fall time becomes relatively short so that a relatively small amount of energy is lost by radiative diffusion through the envelope during the accretion phase. Also, the Kelvin time in the core becomes long, so that relatively little contraction takes place during accretion. Hence, the low-mass stars come into overall hydrostatic equilibrium relatively high on their Hayashi tracks. Protostars with masses in the range 0.5–2 $M_\odot$ become visible with stellar spectra when their radius is about 2 $R_\odot$, which is suspected to be the approximate radius of some of the T Tauri objects. The PMS stars of masses greater than 3 $M_\odot$ would not appear in the H-R diagram as equilibrium objects. Observations show, however, that young stellar objects do exist in this part of the diagram (Strom 1977; Cohen & Kuhi 1976; Cohen 1978) from which we can conclude that the density at the time of initial collapse may be higher than that given by the Jeans criterion at least in some cases. The PMS evolution of a massive star ($\geqslant 10\ M_\odot$) shows that radiation and gas pressure around massive stellar cores limit and ultimately prevent accretion of circumstellar material onto such cores. This suggests that the upper limit to the mass of a star is determined by the properties of cores formed in the centre of massive objects.

Recently, Stahler *et al.* (1980) have proposed a new approach to calculate the early stellar evolution. Their approach is to divide the contracting cloud into regions. In each region certain assumptions are valid which allow the calculations to be simplified. Solutions for infall velocity, pressure, temperature etc. are obtained in each region separately and then the solutions are forced to match at the boundaries between regions. They have ignored the initial phase computations when the entire cloud is in free fall. The end product of the calculations is the results for the end of the accretion phase, when the growing core emerges as an observable star. At this time, the calculations show that the protostar begins to behave according to the 'classical' theory (Hayashi 1961; Henyey *et al.* 1959). The observed positions of the PMS stars on the H-R diagram, as given by Cohen & Kuhi, agree with this classical theory reasonably well. Although the theory by Stahler *et al.*, needs further improvements, its simplicity and the fact that it appears to agree with the observational results of Cohen & Kuhi suggest that a preliminary understanding of PMS evolution may be in sight.

13. Future work

Eventhough the material presented above is not complete, yet, there are areas, in both observations and theory, which need further attention. A complete understanding of the variation of the interstellar reddening law among various young groups is essential in order to estimate the effective temperature, luminosity, mass and age of individual objects that are associated with these young groups. A thorough study of abundance variations and stellar rotation is essential. The role of circumstellar envelopes, gas and dust emission, size, shape, duration of their existence and their role in the PMS evolution need further attention. Finally, a detailed analysis of many of these young groups, both old and new, are urgently needed in order to understand the Galactic evolution.

This review is based on a talk given at the eighth meeting of the Astronomical Society of India held 1982 November at Gorakhpur.

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