Theory of Protostellar Objects

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Abstract

Many problems in the theory of star formation are amenable to a complementary attack in which the analytical approach is used to reduce the governing equations to a form amenable to efficient numerical solution. This strategy has proven very useful in helping to resolve several astrophysical puzzles which arise because the bulk of star formation today is observed to occur, with relatively low efficiency, in giant molecular cloud complexes. How does a cloud of $10^5$-$10^6 \, M_\odot$ know how to form stars of mass $\sim 1 \, M_\odot$? How does the interstellar medium know, to one or two orders of magnitude, that roughly $(\hbar c/G m_H^2)^{3/2}$ hydrogen atoms of mass $m_H$ are needed to yield thermonuclear fusion in a self-gravitating ball of gas? Why have radio astronomers not detected unambiguous evidence for the collapse motions attendant to star formation? Why has a true protostar, the "holy grail" of infrared astronomy, been so hard to find? Why do young stellar objects almost universally exhibit powerful outflows? Why is the geometry for these outflows often bipolar? Why do T Tauri stars have such active chromospheres? In this review we suggest that these puzzles all have a related resolution, in the nature of how gravitational collapse is initiated and terminated in the slowly rotating cores of molecular clouds.

1. Introduction

Following the pioneering work of Hayashi [1] and Larson [2], it is commonly accepted (a) that the birth of a star involves gravitational collapse from an extended cloud of gas and dust, and (b) that a protostar builds up by accretion from the infalling material. This conventional picture presents a considerable challenge for the subject of radiative hydrodynamics because the collapse from interstellar dimensions to stellar ones involves a very large dynamical range, e.g., more than twenty orders of magnitude changes in density. The numerical difficulties associated with following the complete evolution of a protostellar object by pure finite-difference schemes, even in spherical symmetry, led to considerable controversy in the field [3-7]. When the effects of rotation and magnetic fields are included, the numerical problems are compounded [8-17], and to date no theorist has succeeded in rigorously producing stars from interstellar clouds with variations in more than one spatial dimension.

From another point of view, however, it is clear that the existence of large (or small) parameters may offer considerable scope for analytical techniques: simi-
larity solutions [18-21], perturbation techniques [22-24], matched asymptotic expansions [25], etc. Since the finite-difference approach has been amply reviewed elsewhere [26-30], I shall concentrate here on the semi-analytical methods which can be used to reduce the governing equations to a form amenable to quick and accurate numerical solution. Numerical integrations are usually required at the last stage of the analysis if sufficient physical realism is retained to allow meaningful comparisons of the computed results with observational data. Thus, analytical and numerical techniques should be regarded in this field, as in many others, as being complementary rather than being competitive.

2. The Mass Scale of Stars

To begin a discussion of the process of star formation, we should first define what we mean by a (normal) star. The following definition would probably satisfy most astronomers and nuclear physicists: A normal star is a luminous ball of self-gravitating gas which possesses sufficient mass to enable thermonuclear fusion in its central regions. This immediately raises the interesting question: After mass loss, how heavy does an isolated star have to be for fusion reactions to provide a major source of energy? The answer given by stellar evolution theory (see [31] and refs. therein) is, to a level of accuracy sufficient for our purposes,

Deuterium burning only: $0.01M_\odot < M_* < 0.08M_\odot$,
Hydrogen burning also: $0.08M_\odot < M_* < 0.4M_\odot$,
Helium burning also: $0.4M_\odot < M_* < 1.4M_\odot$,
Iron and beyond: $M_* > 1.4M_\odot$.

The value $1.4 M_\odot$ represents, of course, Chandrasekhar's limiting mass for a white dwarf: $0.78(\hbar c/Gm_p^2)^{5/2}m_p$, where $m_p$ is the mass of the proton and the other symbols have their usual meanings. Our understanding of the mass ranges given above rests with the notion that all stars evolve as to try ultimately to produce a white dwarf at its core (for massive stars, this is the pre-supernova state). If this white dwarf has a mass which approaches Chandrasekhar's limiting mass, then it can force the pressures and temperatures in the part of the star just above it to almost arbitrarily high values – values sufficient, indeed, to produce all the chemical species in the periodic table up to iron (by charged particle reactions) and beyond (by neutron capture reactions). On the other hand, if the final white-dwarf core is less than Chandrasekhar's limit, then nuclear burning is terminated at an earlier stage.

The most common stars in our Galaxy and other giant galaxies seem to have masses of $0.5 M_\odot$ or less [32]; thus, within a factor of 3 or so, Chandrasekhar's limit seems to be a natural unit for the mass scale of stars. This central fact can be regarded as the fundamental issue which must be addressed by any viable theory of star formation: How does a galaxy of mass $10^{11}M_\odot$ and size $10^{23}$ cm know how to form objects of mass $10^0 M_\odot$ and size $10^{11}$ cm? Most astronomers today would