Formation of massive stars: a review

Patrick Hennebelle

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Gilles Chabrier, Romain Teyssier
How massive stars form?

- Can we form massive stars in spite of the radiative pressure?
  - 1D: grain solution
  - 2D: flashlight solution
  - 3D: radiative instability solution
  - collision model

- Can we prevent the gas to fragment in many objects?
  - isothermal
  - radiative feedback
  - magnetized non-radiative
  - radiative and magnetized

- Where the gas is coming from?
  - competitive accretion
  - gravitational well based theory
  - tracing the gas in simulations
Thermal Support
Consider a cloud of initial radius $R$

If $\gamma < 4/3$, when $R$ decreases, $\frac{E_{\text{therm}}}{E_{\text{grav}}} = \frac{PV}{GM^2/R} \propto \rho^{\gamma} R^4 \propto R^{4-3\gamma}$ decreases:
$\Rightarrow$ heating/cooling processes

Centrifugal Support and Angular Momentum Conservation

When $R$ decreases, $\frac{E_{\text{rot}}}{E_{\text{grav}}} = \frac{MR^2\omega^2}{GM^2/R} \propto \frac{1}{R}$ increases:
$\Rightarrow$ (magnetic) braking process

$\omega = \frac{R_0^2 \omega_0}{R}$

Magnetic Support and Flux Conservation

When $R$ decreases, $\frac{E_{\text{mag}}}{E_{\text{grav}}} = \frac{B^2R^3}{M^2/R} \propto (\phi/M)^2$ is constant:

Typically one infers $\mu = (M/\phi)/(M/\phi)_c = 1-4$
(Crutcher et al. 1999, 2004)
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The Issue of Radiative Pressure

The coupling occurs through the coupling of the radiation to the dust
=> Treating well the radiative transfer and the micro-physics

Larson & Starrfield 1971 (Kahn 1974):

+ Collapse conditions =>
  Radiation Pressure / Dynamical Pressure = 1

\[ u^2 = 2GM / r \Rightarrow \frac{L}{4\pi r^2 c} \rho \frac{u^2}{u^2} \approx 1.3 \times 10^{-11} \frac{L/L_s}{(M/M_s)^{1/2}} r^{1/2} \]

+ the radius where grains vaporize corresponds to => T=1500 K

+ Estimate of T using simple radiative transfer

=> Radiative Pressure / Dynamical Pressure = \[ 2 \times 10^{-5} \frac{(L/L_s)^{6/5}}{(M/M_s)^{3/5}} \]

leads to M~20 M_s
Wolfire and Cassinelli (1985):
- assume a constant accretion rate of $10^{-3} M_\odot$/an
- consider a classical (Matthis et al. 1977) grains distribution
- explore the influence of a grain deficit

A 100 $M_\odot$ stars cannot form with a standard grain abundance (1/4 is required)

Solution:
weak abundance of grains

1D results by Kuiper et al. (2010)
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The Multi-D approach
(e.g. Yorke & Sonnhalter 2002, Krumholz et al. 2009, Kuiper et al. 2010)

Flux limited diffusion method:

- implicit methods: expensive
- usually assume grey-body except Yorke & Sonnhalter who do multi-frequency
- Kuiper et al. use an hybrid scheme, treating FLD as grey but also include the direct illumination from central stars (multi-wavelength)

\[
\frac{\partial \rho}{\partial t} + \nabla \cdot (\rho \mathbf{v}) = 0,
\]

\[
\frac{\partial (\rho \mathbf{v})}{\partial t} + \nabla \cdot (\rho \mathbf{v} \mathbf{v}) = -\nabla P - \rho \nabla \phi,
\]

\[
\frac{\partial (\rho \mathbf{e})}{\partial t} + \nabla \cdot [(\rho \mathbf{e} + P) \mathbf{v}] = \rho \mathbf{v} \nabla \phi - \kappa_R \rho (4\pi B - cE),
\]

\[
\nabla^2 \phi = 4\pi G \left[ \rho + \sum_n m_n \delta(x - x_n) \right],
\]

\[
\frac{\partial E}{\partial t} - \nabla \cdot \left( \frac{c \lambda}{\kappa_R} \nabla E \right) = \kappa_P \rho (4\pi B - cE) + \sum_n L_n W(x - x_n)
\]

\[
L_n = \frac{G \dot{M} M}{R} + L_*
\]

\[
\kappa_R = 0.1 + 4.4(T_g/350) \text{ cm}^2 \text{ g}^{-1},
\]

\[
\kappa_P = 0.3 + 7.0(T_g/375) \text{ cm}^2 \text{ g}^{-1}.
\]
2D results by Yorke & Sonnhalter (2002)

Solution of the radiative pressure problem: the anisotropy due to the disk, *The flashlight effect* (Yorke & Bodenheimer 99)

-due to centrifugal force, matter piles up in the equatorial plan

-bigger optical depth in the equatorial plan than along the pole

-the photons escape in this direction and the radiative pressure is reduced
Yorke & Sonnhalter
(2D, 64 wavelengths)
+ collapse nearly spherical
+ flattening in the equatorial plan
+ formation of outflows along the pole
+ eventually expansion even along the equatorial plan

Formation of 33 Ms star
(total mass 60).
Yorke & Sonnhalter

Identical but grey treatment

Important differences:

- no outflow

-Final Mass : 20 Ms (total mass 60 Ms).
Results of Kuiper et al. 2010
Importance of solving the dust free regions
(otherwise the radiation is too isotropic and this stops the Flashlight effect)
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3D results by Krumholz et al. 2009
The cavity becomes unstable to Rayleigh-Taylor instability and more gas is accreted.
Results of Kuiper et al. 2011
The instability of the cavity is due to the FLD scheme. The hybrid scheme (FLD+ray tracing) leads to higher radiative pressure and no instability.
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Alternative Scenario (Bonnell et al. 98, 02, 04):
Formation of massive stars by «merging» of low mass stars

The radiative pressure has no influence on the collisions between stars

Bonnell et al. 98 develop a Toy model that takes into account accretion, merging by collision and cluster dynamics.

=> efficiency required 10^8 stars/pc^2
Estimate using direct simulations
(Bonnell & Bate 02,05)
SPH with 1,000,000 of particules, use sink particles
Possible to merge the sinks (distance threshold)

+ Mass of most massive stars
+ Mean mass in the cluster core
+ Mean mass

Mass due to collisions as a function of the final mass
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Fragmentation of a collapsing unmagnetized isothermal Massive core (Dobbs et al. 2005)

The core fragments in many objects therefore limiting the mass of stars.
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Influence of the radiative feedback on a massive core
Krumholz et al. 2007

Radiative feedback reduces the number of fragments simply because it increases the temperature and increases the Jeans mass.

Note however that the initial conditions are very peaked (singular isothermal sphere) and tend to prevent fragmentation.

Girichidis et al. 2010
Influence of the protostellar feedback: further calculations


Bate 2009, 2012

without feedback

with feedback

with feedback

without feedback

Offner et al. 2009
Wind and Radiative feedback

Krumholz et al. 2012

Pure radiation

Wind & radiation

=> Too few low mass objects with pure radiation, winds help to reduce feedback which escape along cavities
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Influence of a weak magnetic field on the fragmentation of low mass cores

\[ \mu = 1000 \text{ (hydro)} \]

\[ \mu = 50 \]

\[ \mu = 20 \]

Low magnetic fields allow disk formation but the disk is stabilized and does not fragment.

H & Teyssier 2008 (see also Machida et al. 2005)
100 M⊙ magnetized, turbulent and dense barotropic core
(other related works: Peters et al. 2010, Seifried et al. 2012)
Turbulence is initially seeded. $\frac{E_{\text{turb}}}{E_{\text{grav}}} \sim 20\%$

In the case of a massive turbulent core, magnetic field reduces, though, do not suppress fragmentation.

$\mu = 120$  $\mu = 5$  $\mu = 2$

H+2011
Impact of the magnetic braking:

J is much reduced as B Increases

=>

Magnetic braking is important even when the flow is significantly turbulent
100 $M_\odot$ magnetized, turbulent and dense barotropic core

Powerful outflows are launched even in turbulent cores
Faster flows appear with weaker fields
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100 M\(_\odot\) turbulent dense core collapse

Eturb/Egrav=20% initially

weak B

strong B


(see also Price & Bate 2009 at larger scales, Myers et al. 2013, 2014)
100 $M_\odot$ turbulent dense core collapse

100 M$_\odot$ turbulent dense core collapse

Trend confirmed with lower resolution runs:

- weak B
- strong B
100 Ms cores with larger column densities
(Myers et al. 2014, Commercon & H 2015)

Similar effects still observed
Magnetized runs have about 2 times less fragments
Ionising radiation tends to push further the gas outwards and may be important/dominant for outflows in massive stars.
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Competitive accretion: individual wells are unimportant
(Zinnecker 1982, Bonnell et al. 2001, Bate et al. 2003,…)

Stellar dominated Potential:
\[ \frac{dN}{dM} \propto M^{-2} \]

Assume \( \rho_{gas} \propto R^{-3/2} \) : typical after rarefaction wave propagates away
(Shu 1977) \( \dot{M}_* = \pi \rho V_{rel} R_{acc}^2 \)

\[ R_{acc} \approx R_{BH} \approx \frac{GM_*}{V_{rel}^2}, \quad V_{rel} \approx \sqrt{GM_{\text{cluster}}/R} \approx R^{-1/2} \]

\[ \Rightarrow \dot{M}_* \propto M_*^2 \] (accretion independent on the position in the cluster)
\[ \Rightarrow dN \propto M_*^{-2} dM_* \] (under reasonable assumptions…)

Mass spectrum from Bonnell et al. (2001)

1000 stars initially of mass 0.1 Ms, 10% of the total Mass

The mass spectrum develops and lead to a Salpeter type Slope
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Theories assuming that individual wells are determinant
(Padoan et al. 97, McKee & Tan 2003, H & Chabrier 2008, Hopkins 2012)

Thermal support
Turbulent compression
+ gravity

Turbulent dispersion
Turbulent compression
+ gravity

Chabrier’s IMF

H & Chabrier 08

\[
N(\tilde{M}) = 2N_0 \frac{1}{R^3} \frac{1}{1 + (2\eta + 1)M^2 R^{2\eta}} \left[ \frac{1 + (1 - \eta)M^2 R^{2\eta}}{(1 + M^2 R^{2\eta})^{3/2}} - \frac{\delta^2 + \sigma^2/2}{(1 + M^2 R^{2\eta})^{1/2}} \right] n' - 3 \frac{\sigma_0^2}{4} \left( \frac{\tilde{R}}{L_i} \right)^{n' - 3}
\times \exp \left\{ - \frac{[\ln(\tilde{M}/R^3)]^2}{2\sigma^2} \right\} \frac{\exp(-\sigma^2/8)}{\sqrt{2\pi\sigma}}
\]
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A hierarchy of wells:

Direct mapping between the wells and the stars?
Exchange between the wells?
Likely both! But how much?
Which mechanism is at play in gravo-turbulent simulations?

Competitive accretion or core formation?

Smith et al. have run SPH simulations with gravity and sink particles.

They identify cores and look at the correlation between the core masses and the sink masses.

The correlation is very good initially (few freefall times) and becomes progressively less good.

=> This is compatible with the core mass function being able to produce a reasonable IMF (Chabrier & H 2010).

Until how many freefall times are the cores accreting?

At the end of the simulation

The most massive stars is more massive than the mass contains in its initial well.
Where this mass comes from?
Where the gas comes from?

Green gas eventually falls in the massive star.

Smith et al. 2009
A hierarchy of wells:

Direct mapping between the wells and the stars?
Exchange between the wells?
Likely both! But how much?
Conclusions

Impact of radiative transfer and magnetic field are obviously drastic in:
- regulating the mass accretion (but not stopping it)
- limiting the fragmentation

**Combination** of magnetic field and radiative transfer is more than their mere juxtaposition.

Where the mass is coming from? still unclear.
=> **Accretion is not a sufficiently clear statement.**

The salient questions are:
- is accretion determined by the present mass object (compet. accret.)?
- is accretion determined by the initial well hierarchy?

The feedback from massive stars is very non-linear. But predicting masses with an accuracy better than ~2 is a huge challenge.
=> a huge multi-scale, multi-physics problems
100 M☉ magnetized, turbulent and dense barotropic core

Powerful outflows are launched even in turbulent cores
Faster flows appear with weaker fields
Growth of the toroidal magnetic field within the disk

Importance of $V_a/C_s$ for various $\mu$ and various times

$\Rightarrow$ Compatible with the assumption that the toroidal field stabilizes the disk.
The final mass of the stars is about 35-37 Ms.
The history depends on the core mass.