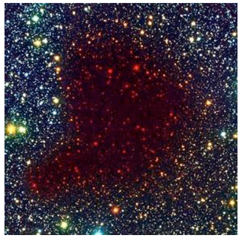
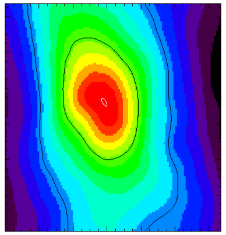
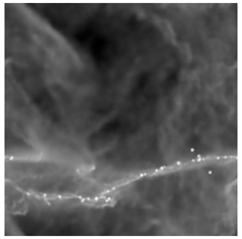
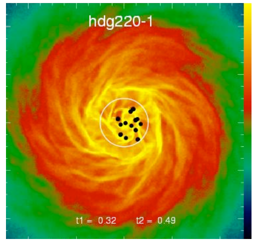
Molecular cloud dynamics and star formation





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Schedule

11h30--12h30	C6.1.	Formation of molecular clouds
16h00--17h00	C6.2.	Origin and statistical characteristics of ISM turbulence
17h00--18h00	C6.3.	Star (cluster) formation in molecular clouds
18h00--19h00	C6.4.	Stellar initial mass function

Lecture 3 + 4: star (cluster) formation and the IMF

- ingrediences of star (cluster) formation
  - dynamics in gas and stars
  - importance of thermydynamics
  - effects of magnetic fields
  - radiative feedback
- different IMF models
  - comments on low-mass end
  - comments on high-mass end

theory

Early dynamical theory

- *Jeans (1902):* Interplay between self-gravity and thermal pressure
  - stability of homogeneous spherical density enhancements against gravitational collapse
  - dispersion relation:

$$\omega^2 = c_s^2 k^2 - 4\pi G \rho_0$$

  - instability when $\omega^2 < 0$
  - minimal mass: $M_J = \frac{1}{6}\pi^{-5/2} G^{-3/2} \rho_0^{-1/2} c_s^3 \propto \rho_0^{-1/2} T^{-3/2}$

Sir James Jeans, 1877 - 1946

(full detail in Mac Low & Klessen, 2004, Rev. Mod. Phys., 76, 125-194)

First approach to turbulence

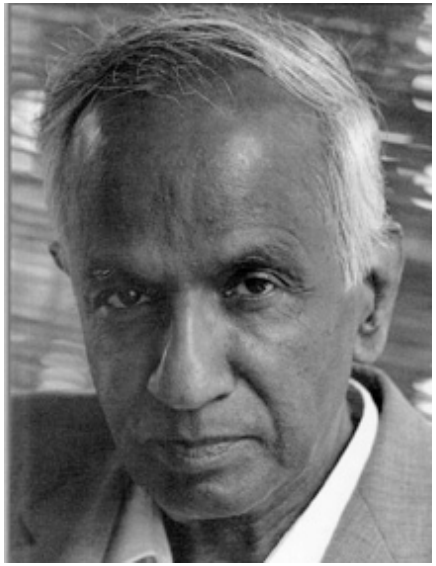
- *von Weizsäcker (1943, 1951)* and *Chandrasekhar (1951):* concept of **MICROTURBULENCE**
  - BASIC ASSUMPTION: separation of scales between dynamics and turbulence

    $\ell_{turb} \ll \ell_{dyn}$
  - then turbulent velocity dispersion contributes to effective soundspeed:

$$c_c^2 \mapsto c_c^2 + \sigma_{rms}^2$$

  - → Larger effective Jeans masses → more stability
  - BUT: *(1) turbulence depends on k:* $\sigma_{rms}^2(k)$

    (2) supersonic turbulence → $\sigma_{rms}^2(k) >> c_s^2$ *usually*



S. Chandrasekhar, 1910 - 1995

(full detail in Mac Low & Klessen, 2004, Rev. Mod. Phys., 76, 125-194)

Problems of early dynamical theory

- Molecular clouds are *highly Jeans-unstable*
  Yet, they do *NOT* form stars at high rate
  and with high efficiency.
  (the observed global SFE in molecular clouds is *~5%*)
  → *something prevents large-scale collapse.*
- All throughout the early 1990's, molecular clouds had been thought to be long-lived quasi-equilitrium entities.
- Molecular clouds are *magnetized.*

(full detail in Mac Low & Klessen, 2004, Rev. Mod. Phys., 76, 125-194)

Magnetic star formation

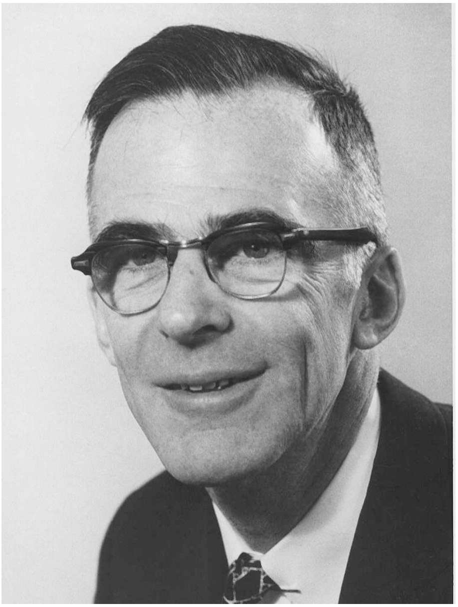
- *Mestel & Spitzer (1956):* Magnetic fields can prevent collapse*!!!*
  - Critical mass for gravitational collapse in presence of B-field

$$M_{cr} = \frac{5^{3/2}}{48\pi^2} \frac{B^3}{G^{3/2}\rho^2}$$

  - Critical mass-to-flux ratio (Mouschovias & Spitzer 1976)

$$\left[\frac{M}{\Phi}\right]_{cr} = \frac{\zeta}{3\pi}\left[\frac{5}{G}\right]^{1/2}$$

  - Ambipolar diffusion can initiate collapse

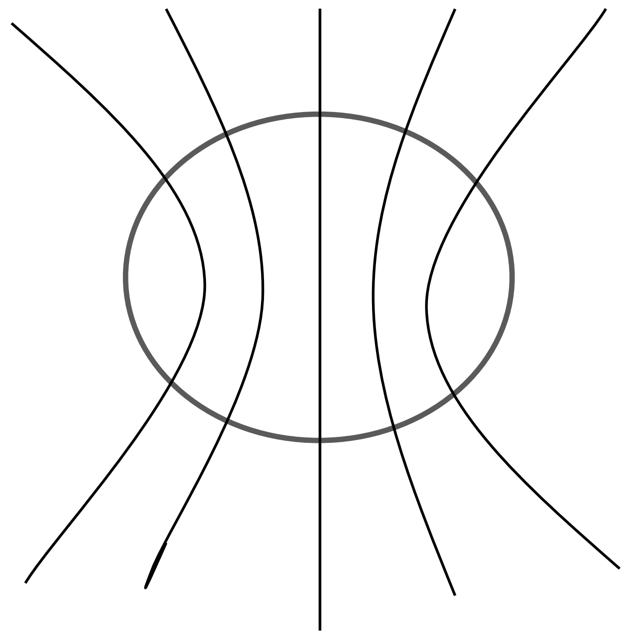


Lyman Spitzer, Jr., 1914 - 1997

(full detail in Mac Low & Klessen, 2004, Rev. Mod. Phys., 76, 125-194)

The "standard theory" of star formation:

- BASIC ASSUMPTION: Stars form from magnetically highly subcritical cores
- Ambipolar diffusion slowly increases (M/Φ): $\tau_{AD} \approx 10\tau_{ff}$
- Once $(M/\Phi) > (M/\Phi)_{crit}$: dynamical collapse of SIS
  - Shu (1977) collapse solution
  - $dM/dt = 0.975\ c_s^3/G$ = const.
- Was (in principle) only intended for isolated, low-mass stars



Problems of magnetic SF

- Observed B-fields are weak, at most marginally critical (Crutcher 1999, Bourke et al. 2001)
- Magnetic fields cannot prevent decay of turbulence
  (Mac Low et al. 1998, Stone et al. 1998, Padoan & Nordlund 1999)
- Structure of prestellar cores
  (Bacman et al. 2000, e.g. Barnard 68 from Alves et al. 2001)
- Strongly time varying dM/dt
  (e.g. Hendriksen et al. 1997, André et al. 2000)
- More extended infall motions than predicted by the standard model
  (Williams & Myers 2000, Myers et al. 2000)

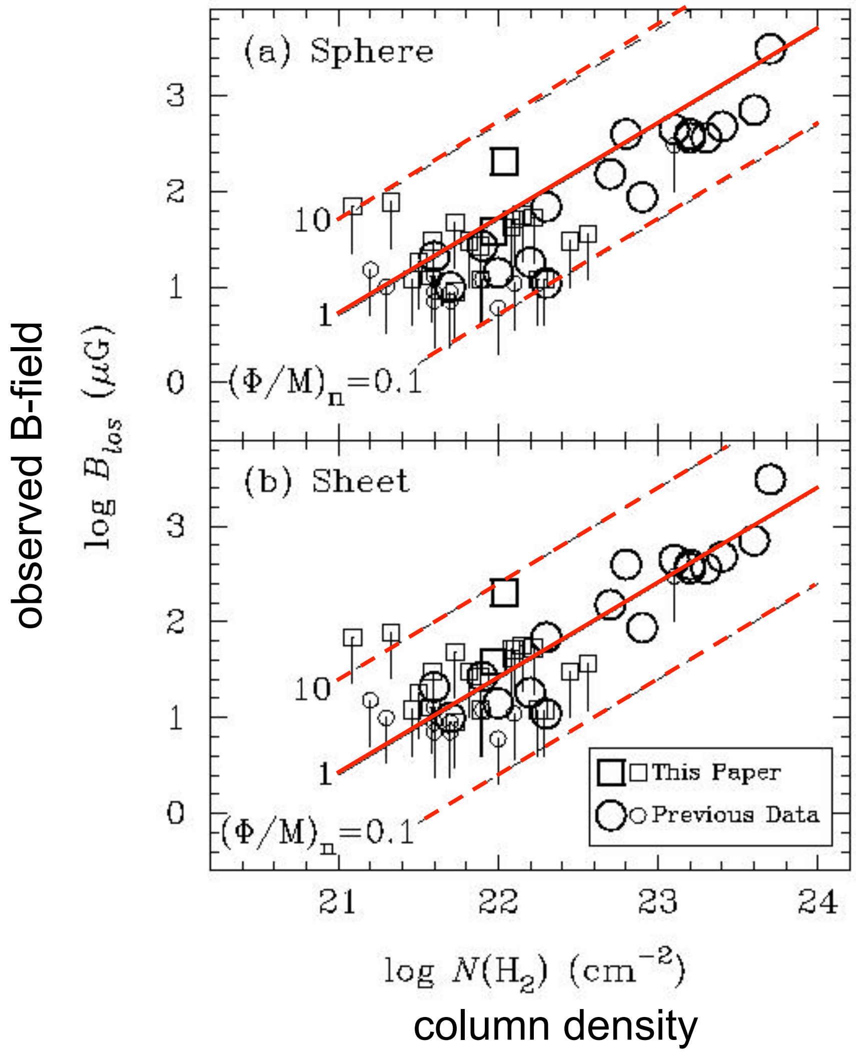
Observed B-fields are weak

B versus ***N(H₂)*** from *Zeeman measurements*.
(from Bourke et al. 2001)

→ cloud cores are magnetically supercritical***!!!***

$(\Phi/M)_n > 1$ no collapse

$(\Phi/M)_n < 1$ collapse



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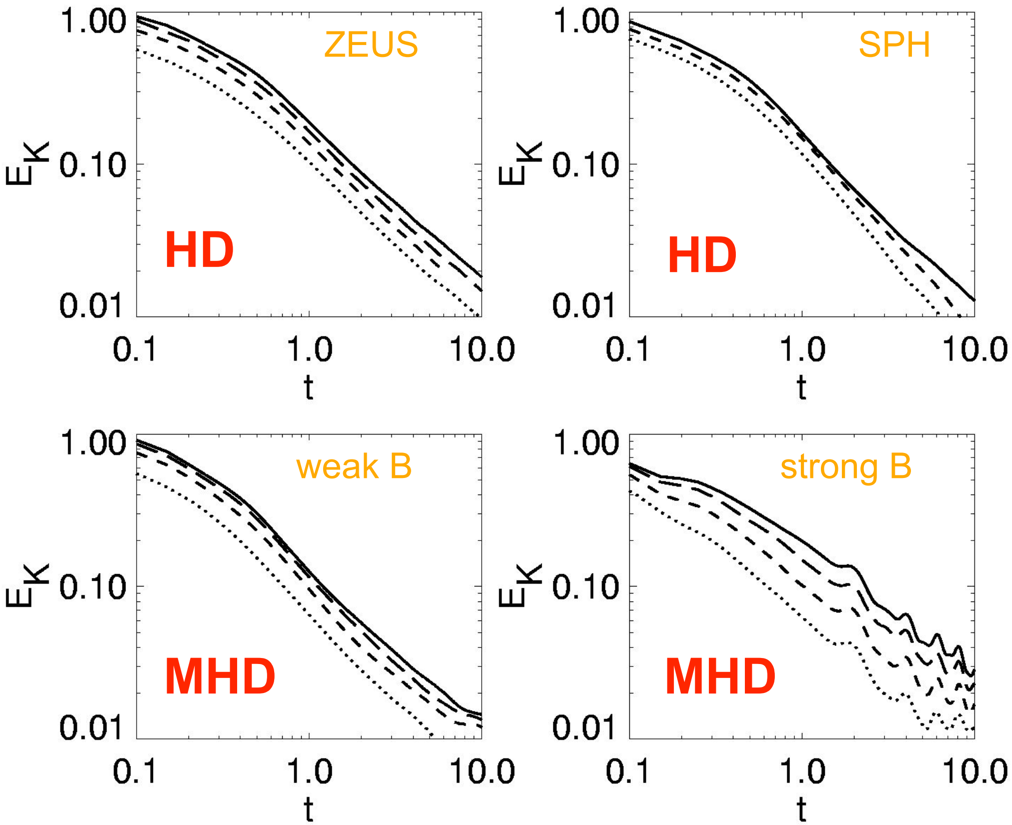
Molecular cloud dynamics

- **Timescale problem:** Turbulence *decays* on timescales *comparable to the free-fall time* τ_{ff} ($E \propto t^{-\eta}$ with $\eta \approx 1$).

  (Mac Low et al. 1998,
  Stone et al. 1998,
  Padoan & Nordlund 1999)

- Magnetic fields (static or wave-like) *cannot* prevent loss of energy.



(Mac Low, Klessen, Burkert, & Smith, 1998, PRL)

Problems of magnetic SF

- As many prestellar cores as protostellar cores in SF regions (e.g. André et al 2002)
- Molecular cloud clumps seem to be chemically young
  (Bergin & Langer 1997, Pratap et al 1997, Aikawa et al 2001)
- Stellar age distribution small ($\tau_{ff} << \tau_{AD}$)
  (Ballesteros-Paredes et al. 1999, Elmegreen 2000, Hartmann 2001)
- Strong theoretical criticism of the SIS as starting condition for gravitational collapse
  (e.g. Whitworth et al 1996, Nakano 1998, as summarized in Klessen & Mac Low 2004)
- Most stars form as binaries

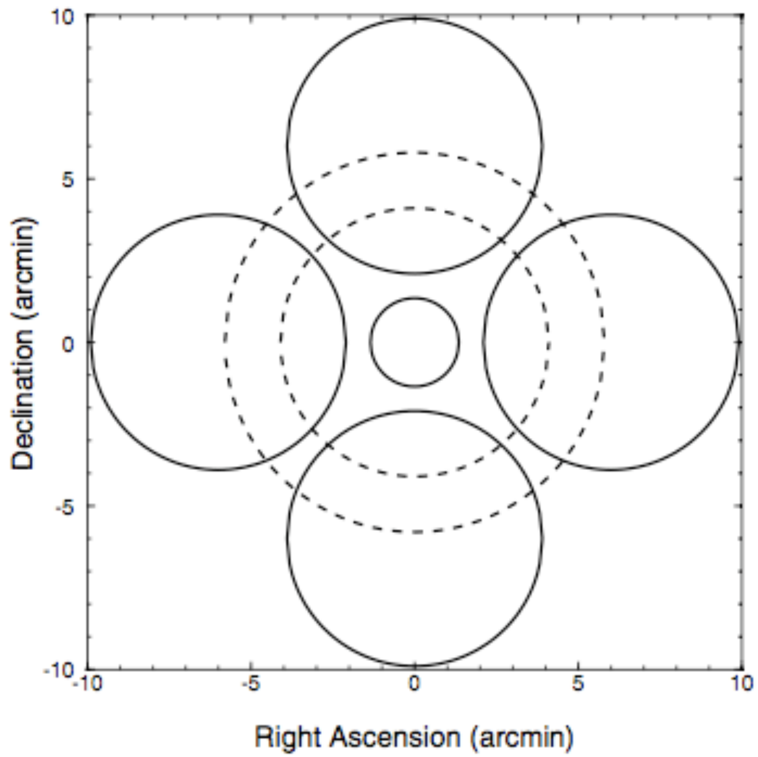
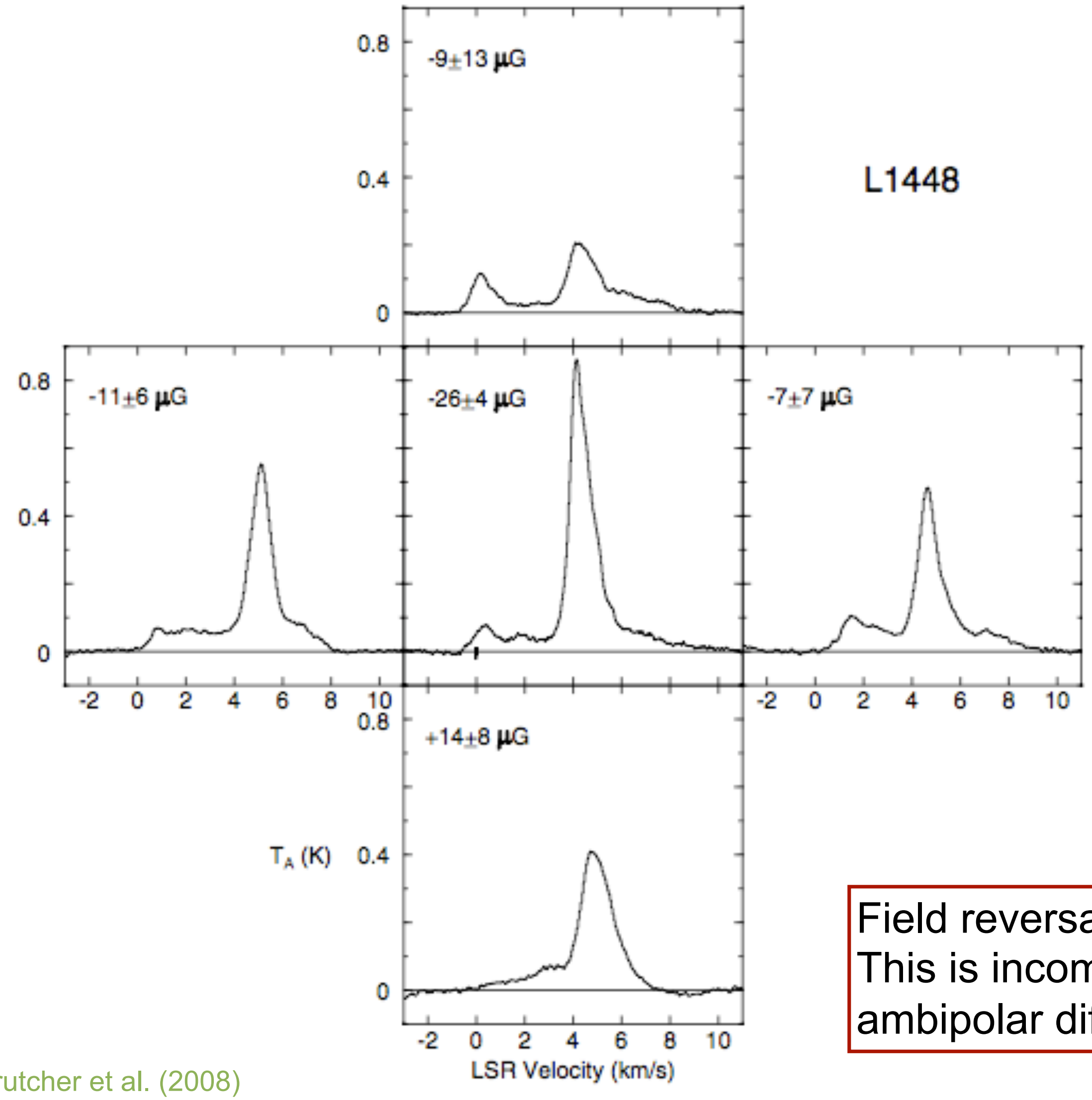


Fig. 1.— The Arecibo telescope primary beam (small circle centered at 0,0) and the four GBT telescope primary beams (large circles centered 6′ north, south, east, and west of 0,0. The dotted circles show the first sidelobe of the Arecibo telescope beam. All circles are at the half-power points.

Crutcher et al. (2008)



Field reversal in the outer parts.
This is incompatible with “standard” ambipolar diffusion theory!

Crutcher et al. (2008)

Fig. 2.— OH 1667 MHz spectra toward the core of L1448CO obtained with the Arecibo telescope (center panel) and toward each of the envelope positions 6′ north, south, east, and west of the core, obtained with the GBT. In the upper left of each panel is the inferred B_{LOS} and its 1σ uncertainty at that position. A negative B_{LOS} means the magnetic field points toward the observer, and vice versa for a positive B_{LOS}.

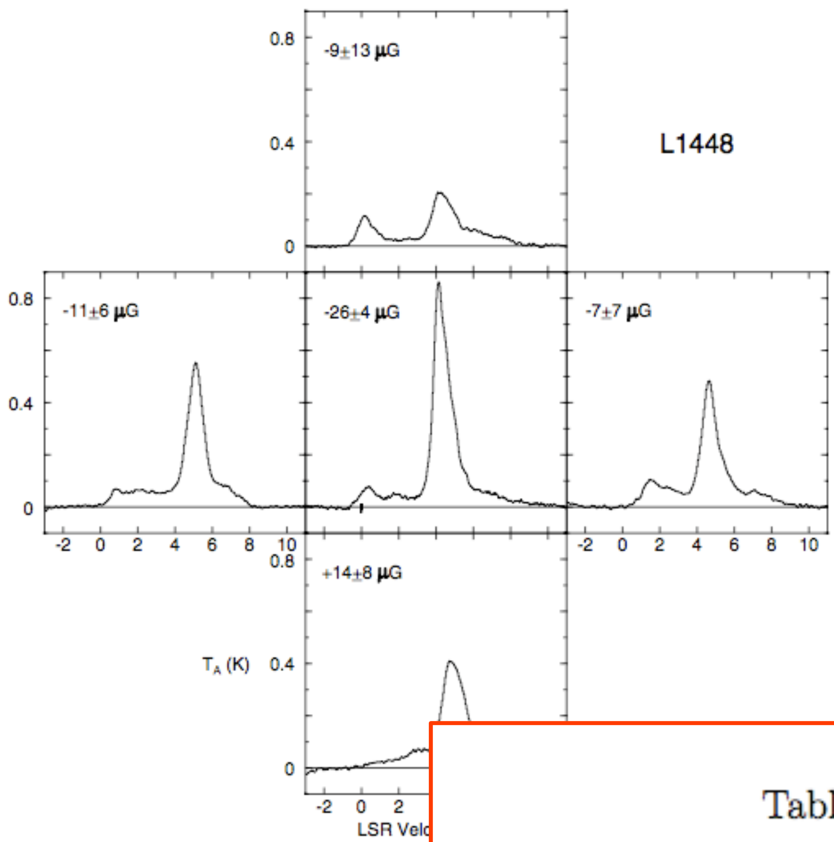


Fig. 2.— OH 1667 MHz spectra toward the
telescope (center panel) and toward each of the
west of the core, obtained with the GBT. In the
and its 1σ uncertainty at that position. A neg
toward the observer, and vice versa for a posit

example: L1448

Table 2. Relative Mass/Flux

Cloud	$\mathcal{R}$	$\mathcal{R}'$	Probability $\mathcal{R}$ or $\mathcal{R}' > 1$
L1448CO	0.02 ± 0.36	0.07 ± 0.34	0.005
B217-2	0.15 ± 0.43	0.19 ± 0.41	0.05
L1544	0.42 ± 0.46	0.46 ± 0.43	0.11
B1	0.41 ± 0.20	0.44 ± 0.19	0.010

Crutcher et al. (2008)

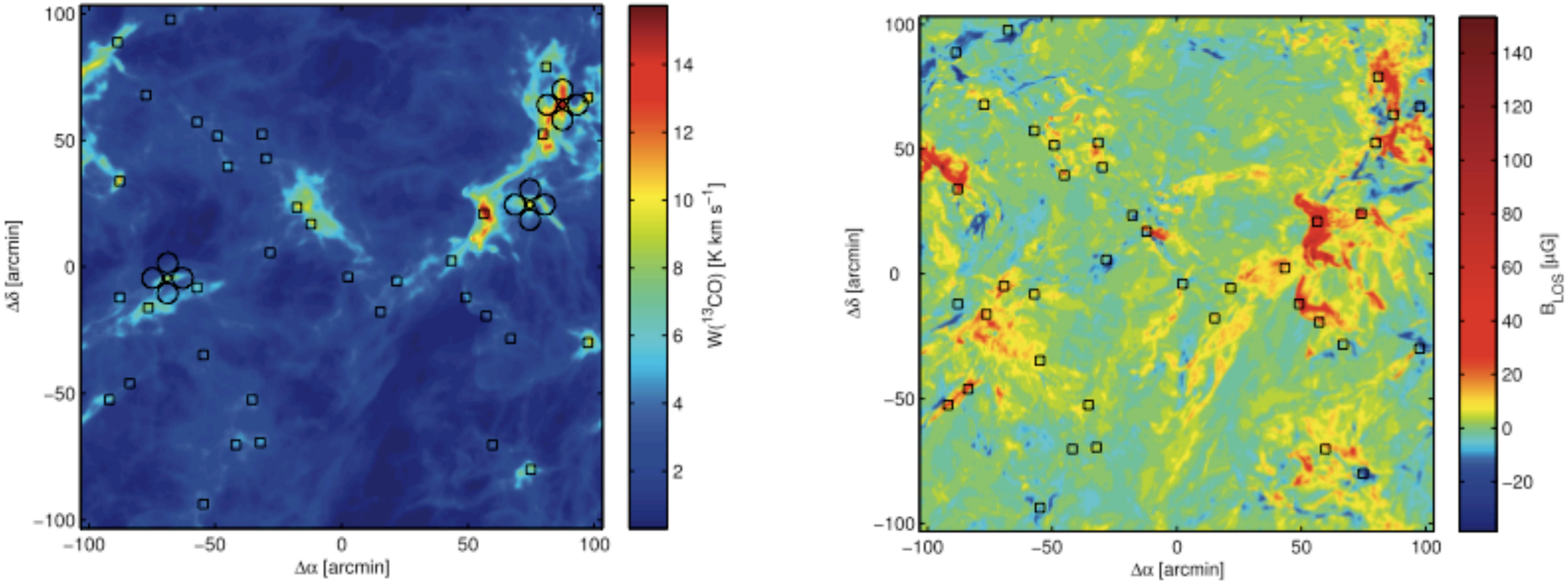


FIG. 1.—*Left*: Simulated ^{13}CO (1–0) map of the model in the *z*-axis direction. The locations of the cloud cores are shown with squares. The circles indicate the locations of telescope beams used in the synthetic observations of three cores. *Right*: Line-of-sight magnetic field strength as calculated from Zeeman splitting.

Lunttila et al. (2008)

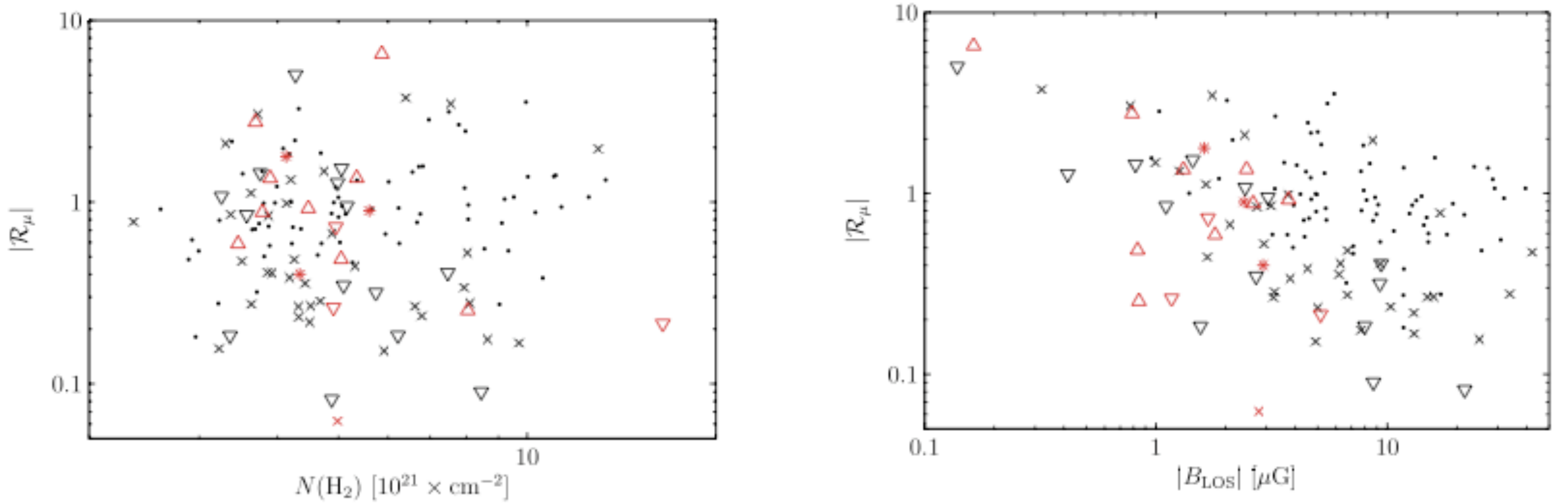


FIG. 3.—*Left*: Relative mass-to-flux ratio for the selected cores as a function of column density. Red symbols indicate the cores with $\mathcal{R}_\mu < 0$. Dots, crosses, triangles pointing down, triangles pointing up, and asterisks denote zero, one, two, three, or four field reversals in the envelopes relative to the core center. *Right*: Relative mass-to-flux ratio as a function of inferred magnetic field strength in the central beam. The symbols have the same meaning as in the left panel.

Lunttila et al. (2008)

gravoturbulent star formation

- idea:

> **Star formation is controlled by interplay between gravity and supersonic turbulence!**

- dual role of turbulence:
  - *stability on large scales*
  - *initiating collapse on small scales*

(e.g., Larson, 2003, Rep. Prog. Phys, 66, 1651;
or Mac Low & Klessen, 2004, Rev. Mod. Phys., 76, 125)

gravoturbulent star formation

- idea:

Star formation is controlled
by interplay between
gravity and
supersonic turbulence!

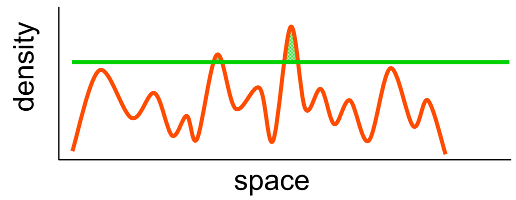
- validity:

This hold on *all* scales and applies to build-up of stars and star clusters within molecular clouds as well as to the formation of molecular clouds in galactic disk.

(e.g., Larson, 2003, Rep. Prog. Phys, 66, 1651;
or Mac Low & Klessen, 2004, Rev. Mod. Phys., 76, 125)

gravoturbulent star formation

- interstellar gas is highly *inhomogeneous*
  - *gravitational instability*
  - *thermal instability*
  - *turbulent compression* (in shocks $\delta\rho/\rho \propto \mathcal{M}^2$; in atomic gas: $\mathcal{M} \approx 1...3$)



- cold *molecular clouds* can form rapidly in high-density regions at *stagnation points* of *convergent large-scale flows*
  - chemical *phase transition*: atomic → molecular
  - process is *modulated* by large-scale *dynamics* in the galaxy
- inside *cold clouds:* turbulence is highly supersonic ($\mathcal{M} \approx 1...20$)
  → *turbulence* creates large density contrast,
  gravity selects for collapse

  ⟶ ***GRAVOTUBULENT FRAGMENTATION***

- *turbulent cascade:* local compression *within* a cloud provokes collapse → formation of individual *stars* and *star clusters*

different statistical approaches

- there are different quantitative IMF based on turbulence
  - Padoan & Nordlund (2002, 2007)
  - Hennebelle & Chabrier (2008, 2009)
  - both relate the mass spectrum to statistical characteristics of the turbulent velocity fields

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ANALYTICAL THEORY FOR THE INITIAL MASS FUNCTION: CO CLUMPS AND PRESTELLAR CORES

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The Astrophysical Journal, 684:395–410, 2008 September 1

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The Astrophysical Journal, 702:1428–1442, 2009 September 10
doi:10.1088/0004-637X/702/2/1428

ANALYTICAL THEORY FOR THE INITIAL MASS FUNCTION. II. PROPERTIES OF THE FLOW

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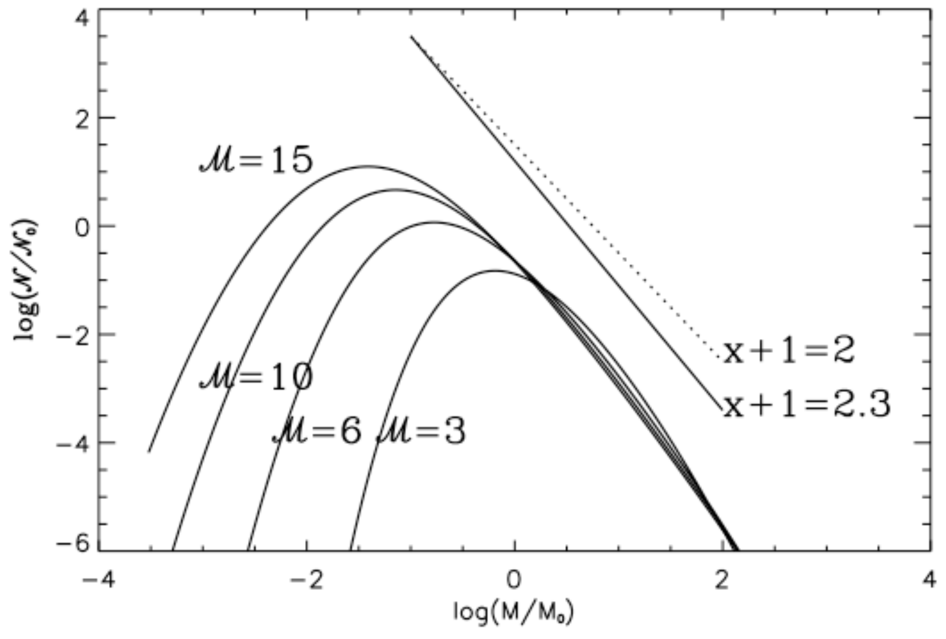
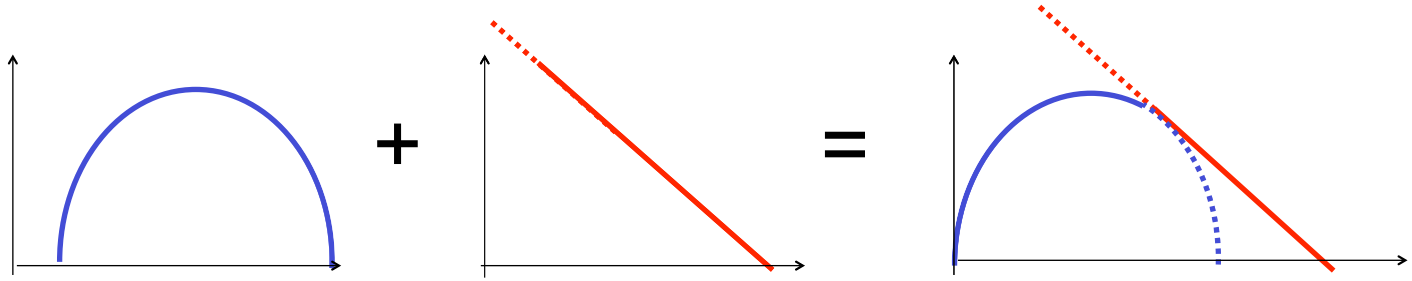


FIG. 2.—Mass spectrum for $\mathcal{M}_*^2 = 2$ and various values of $\mathcal{M}$.

different statistical approaches

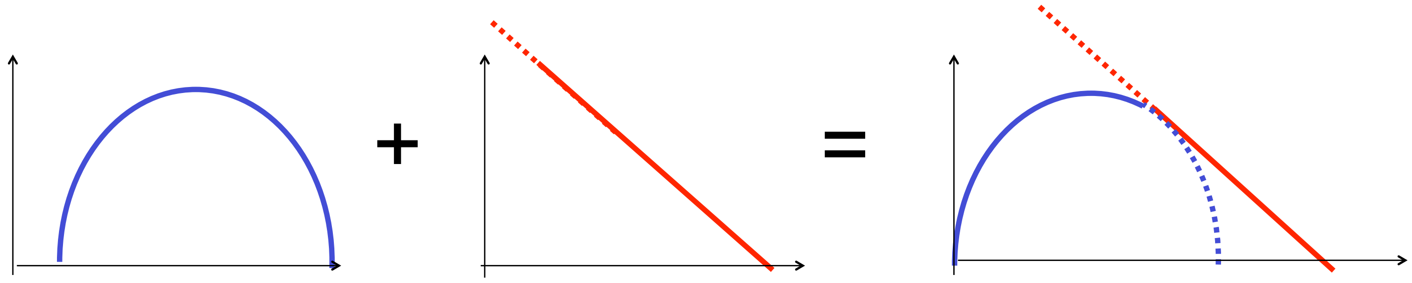
- there are different quantitative IMF based on turbulence
  - Padoan & Nordlund (2002, 2007)
  - Hennebelle & Chabrier (2008, 2009)
  - both relate the mass spectrum to statistical characteristics of the turbulent velocity fields
- there are alternative approaches
  - IMF as closest packing problem / *sampling* problem in *fractal* clouds (Larson 1992, 1995, Elmegreen 1997ab, 2000ab, 2002)
  - IMF as purely *statistical* problem (Larson 1973, Zinnecker 1984, 1990, Adams & Fatuzzo 1996)
  - IMF from (proto)stellar *feedback* (Silk 1995, Adams & Fatuzzo 1996)
  - IMF from competitive *coagulation* (Murray & Lin 1995, Bonnell et al. 2001ab, etc.)

caveat: everybody gets the IMF!



- combine scale free process → *POWER LAW BEHAVIOR*
  - turbulence (Padoan & Nordlund 2002, Hennebelle & Chabrier 2008)
  - gravity in dense clusters (Bonnell & Bate 2006, Klessen 2001)
  - universality: dust-induced EOS kink insensitive to radiation field (Elmegreen et al. 2008)
- with highly stochastic processes → central limit theorem
  → *GAUSSIAN DISTRIBUTION*
  - basically mean thermal Jeans length (or feedback)
  - universality: insensitive to metallicity (Clark et al. 2009, submitted)

caveat: everybody gets the IMF!

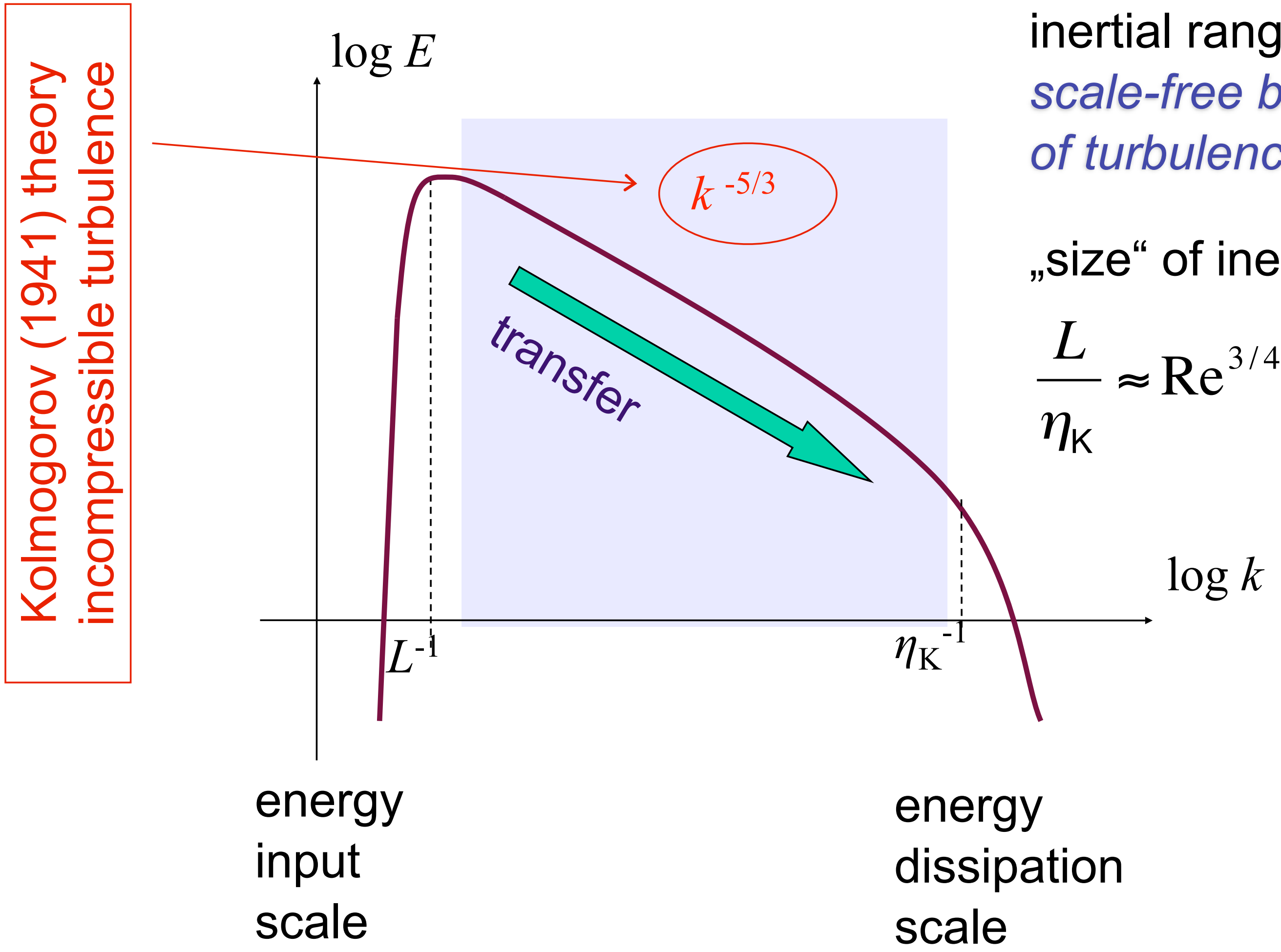


"everyone" gets the right IMF
→ better look for secondary indicators

- *stellar multiplicity*
- protostellar *spin* (including disk)
- *spatial distribution* + *kinematics* in young clusters
- *magnetic field strength* and *orientation*

dynamical approach

Turbulent cascade

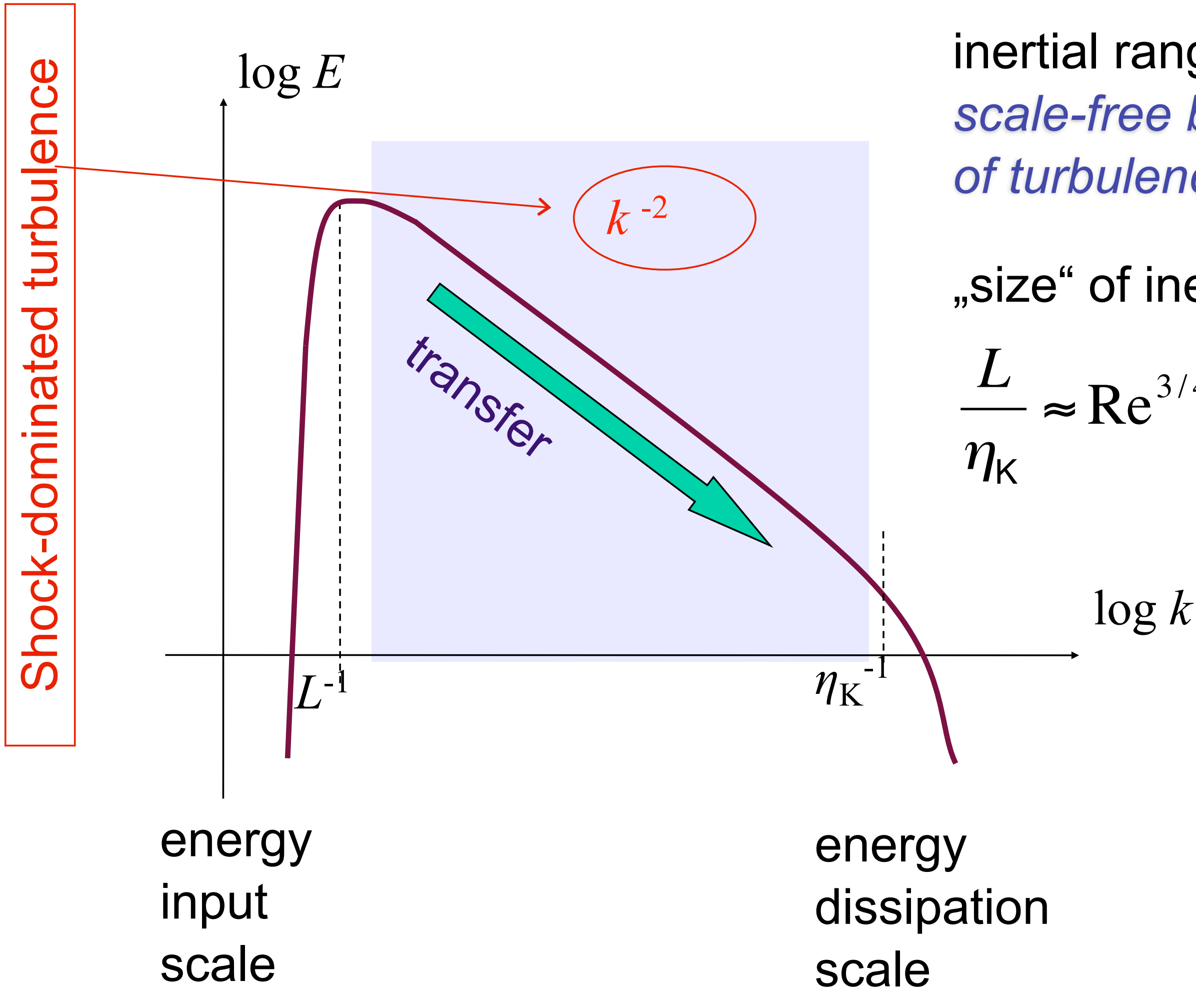


inertial range:
scale-free behavior of turbulence

„size“ of inertial range:

$$\frac{L}{\eta_K} \approx \mathrm{Re}^{3/4}$$

Turbulent cascade



inertial range:
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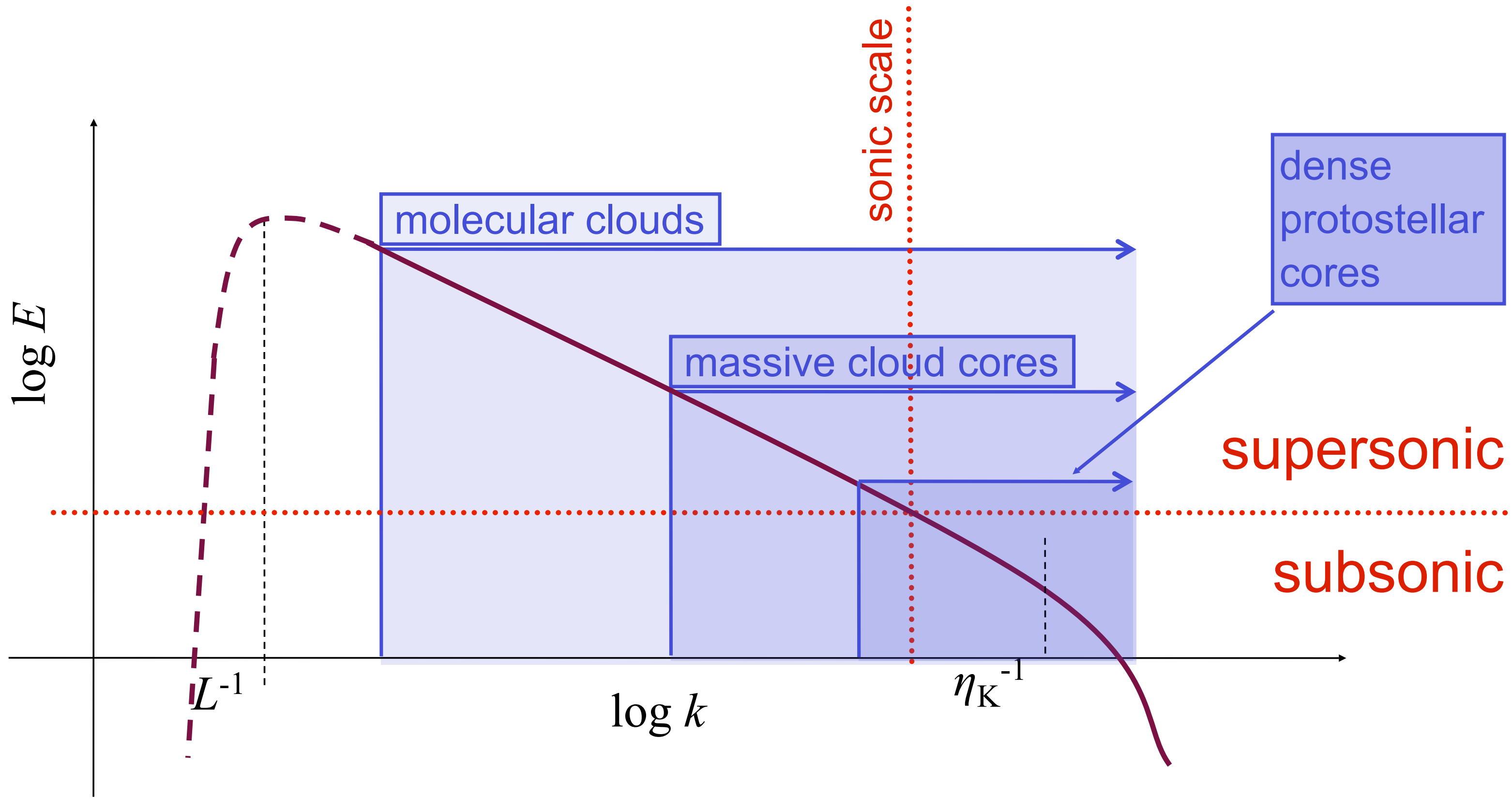
$$\frac{L}{\eta_K} \approx \mathrm{Re}^{3/4}$$

Turbulent cascade in ISM

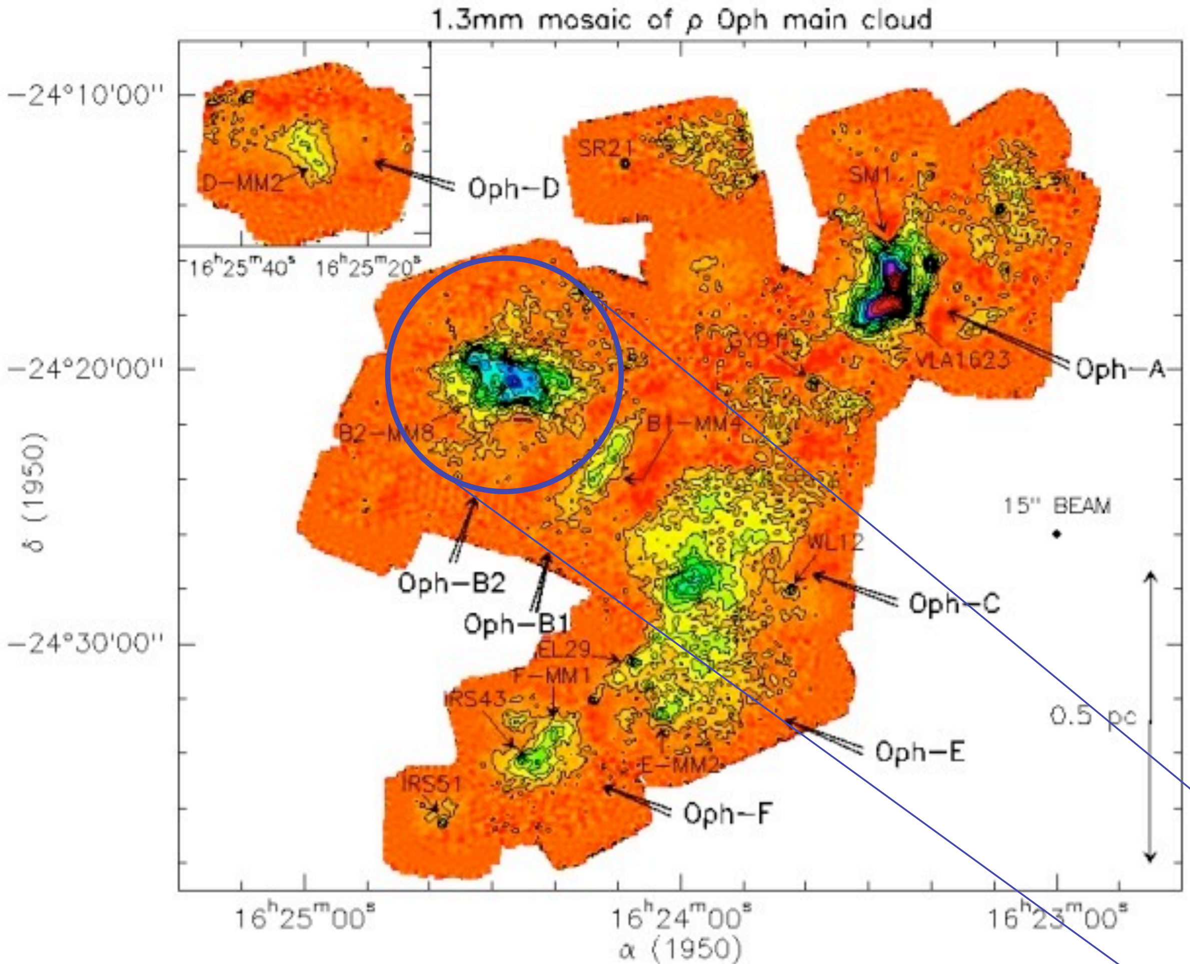


energy source & scale *NOT known* (supernovae, winds, spiral density waves?)

σ_{rms} << 1 km/s
$M_{rms} \leq 1$
L ≈ 0.1 pc

dissipation scale not known (ambipolar diffusion, molecular diffusion?)

Density structure of MC's



(Motte, André, & Neri 1998)

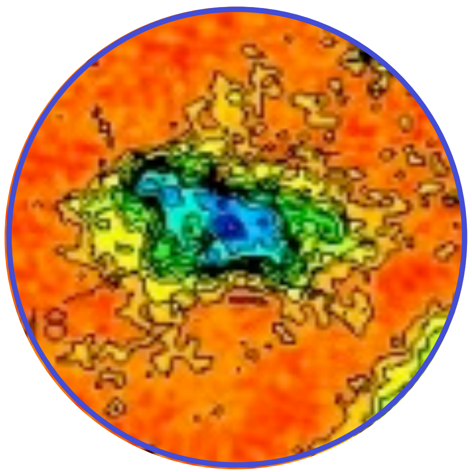
molecular clouds are highly inhomogeneous

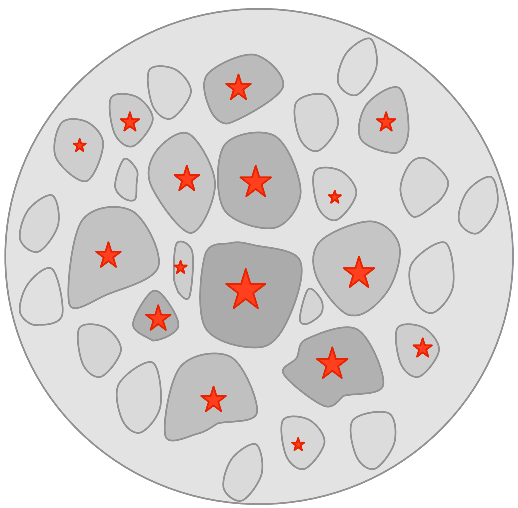
stars form in the densest and coldest parts of the cloud

ρ-Ophiuchus cloud seen in dust emission

let's focus on a cloud core like this one

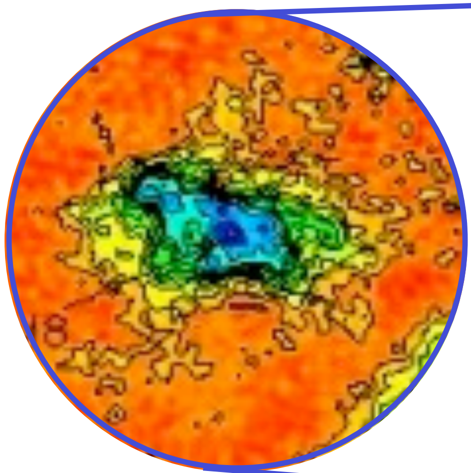
Evolution of cloud cores

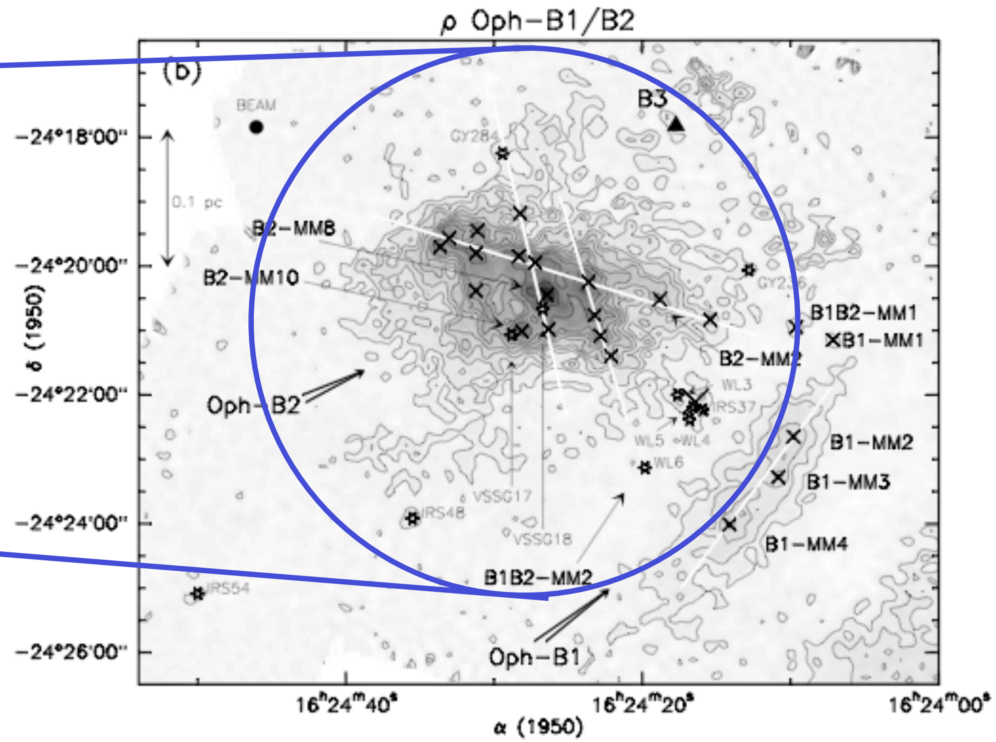


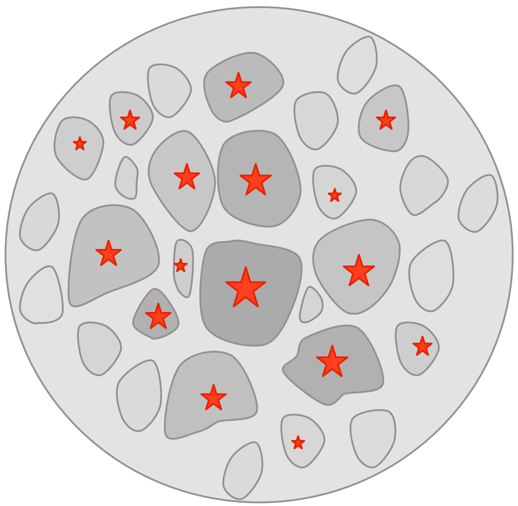


- How does this core evolve?
  Does it form one single massive star or cluster with mass distribution?
- Turbulent cascade „goes through" cloud core
  --> NO *scale separation* possible
  --> NO *effective sound speed*
- Turbulence is supersonic!
  --> produces strong density contrasts:
  $\delta\rho/\rho \approx M^2$
  --> with typical $M \approx 10$ --> $\delta\rho/\rho \approx 100$!
- many of the shock-generated fluctuations are Jeans unstable and go into collapse
- --> expectation: *core breaks up and forms a cluster of stars*

Evolution of cloud cores





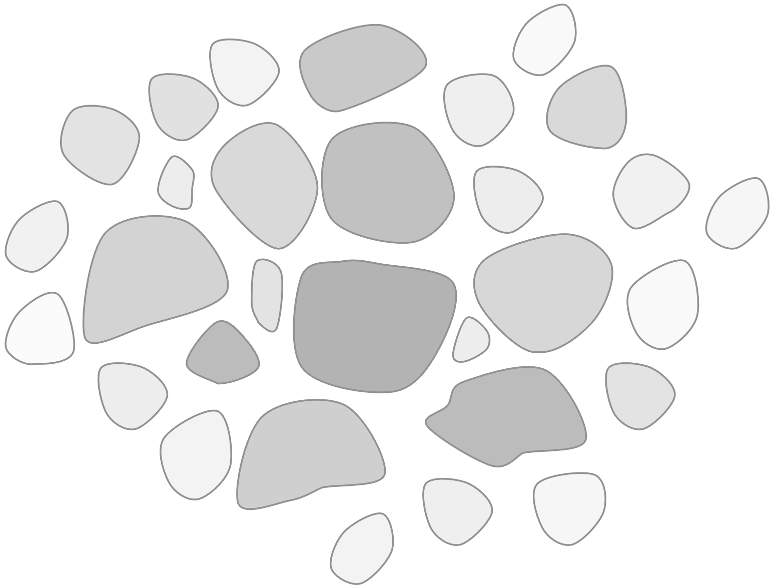


indeed ρ-Oph B1/2 contains several cores ("starless" cores are denoted by ×, cores with embedded protostars by ☆)

(Motte, André, & Neri 1998)

Formation and evolution of cores

What happens to distribution of cloud cores?



Two exteme cases:

(1) turbulence dominates energy budget:

$\alpha = E_{kin}/|E_{pot}| > 1$

--> individual cores do *not* interact

--> *collapse* of *individual* cores dominates *stellar mass growth*

--> *loose cluster of low-mass stars*

(2) turbulence decays, i.e. gravity dominates:

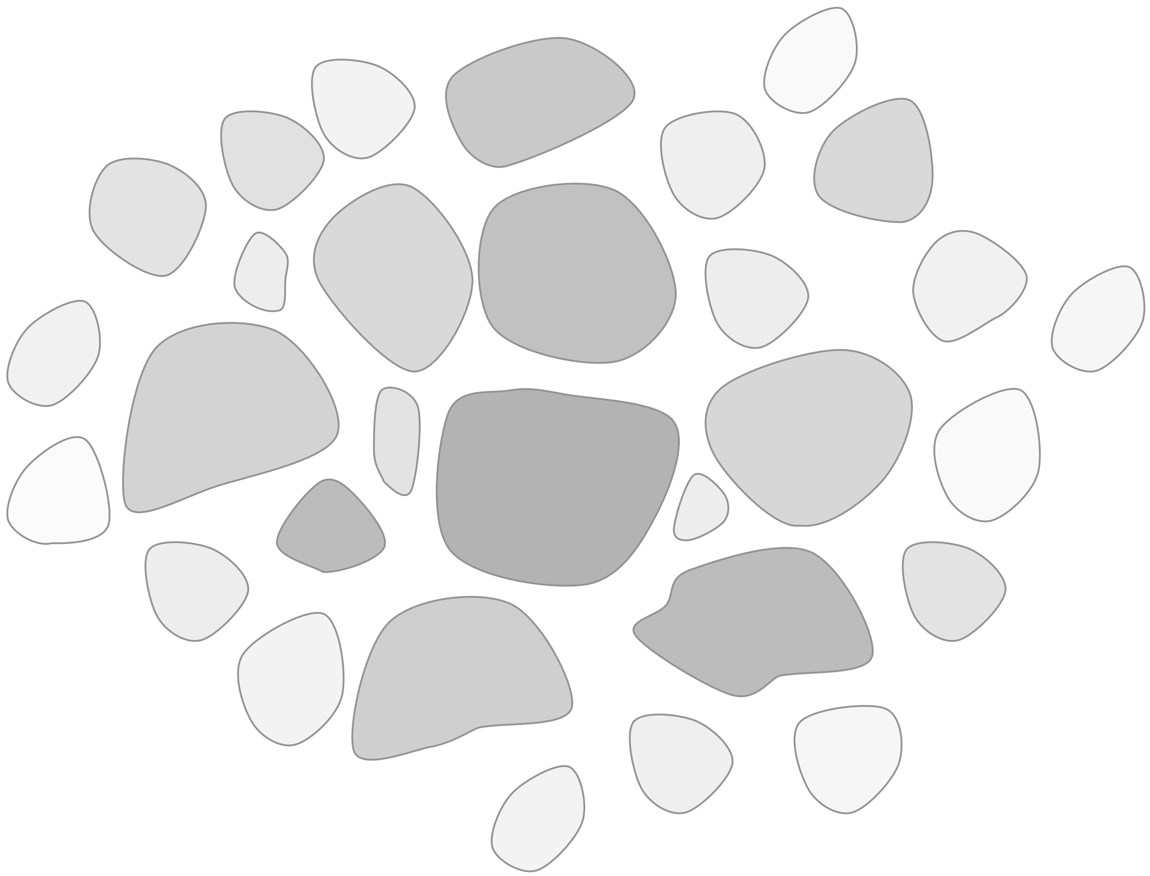
$\alpha = E_{kin}/|E_{pot}| < 1$

--> *global contraction*

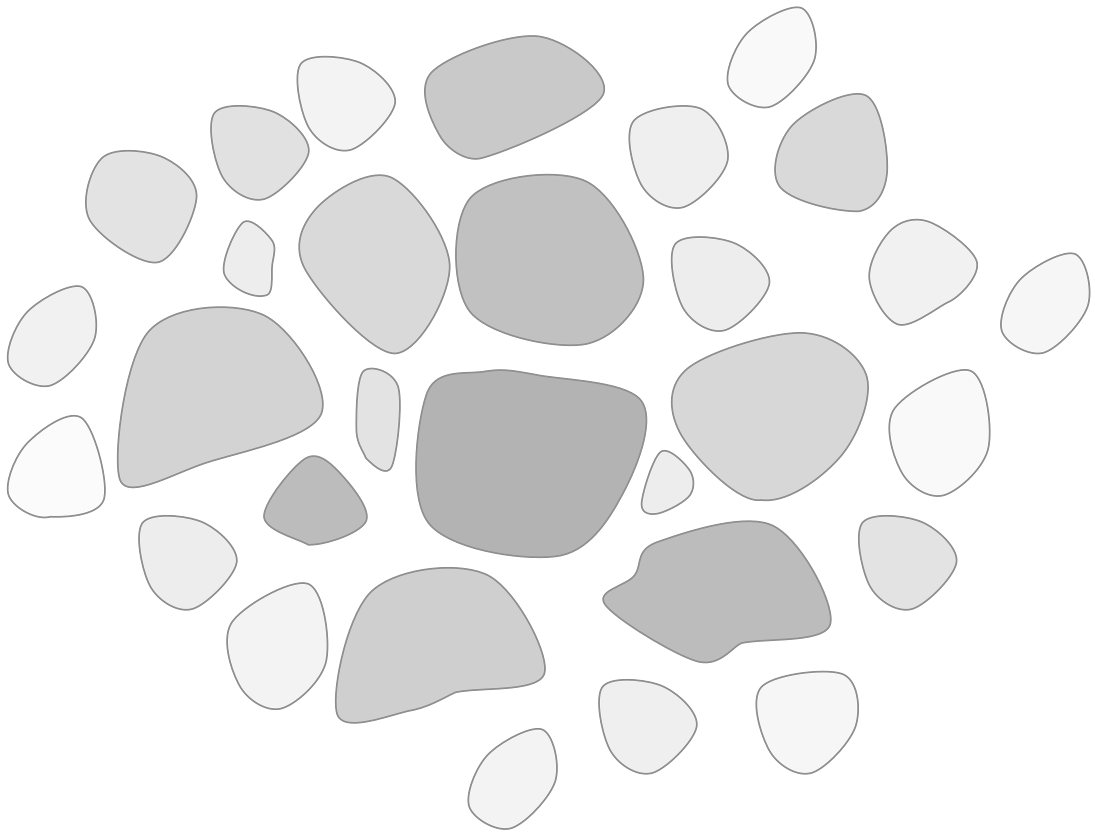
--> core do *interact* while collapsing

--> *competition* influences *mass growth*

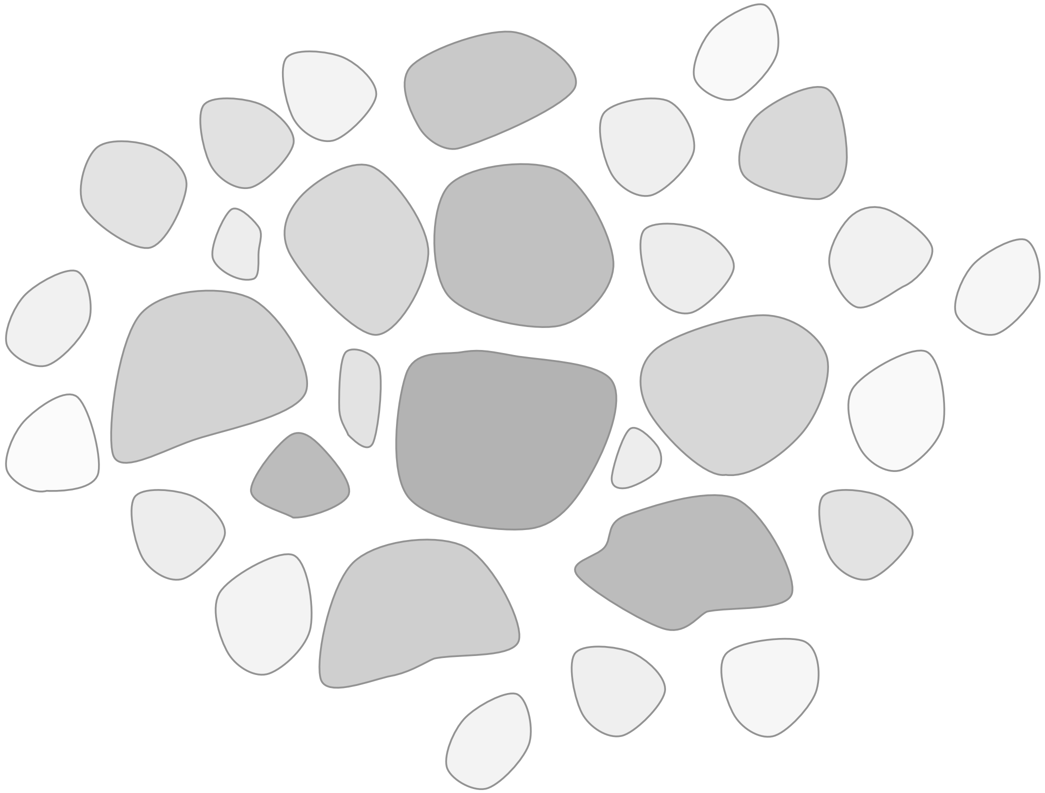
--> *dense cluster with high-mass stars*



turbulence creates a hierarchy of clumps



as turbulence decays locally, contraction sets in



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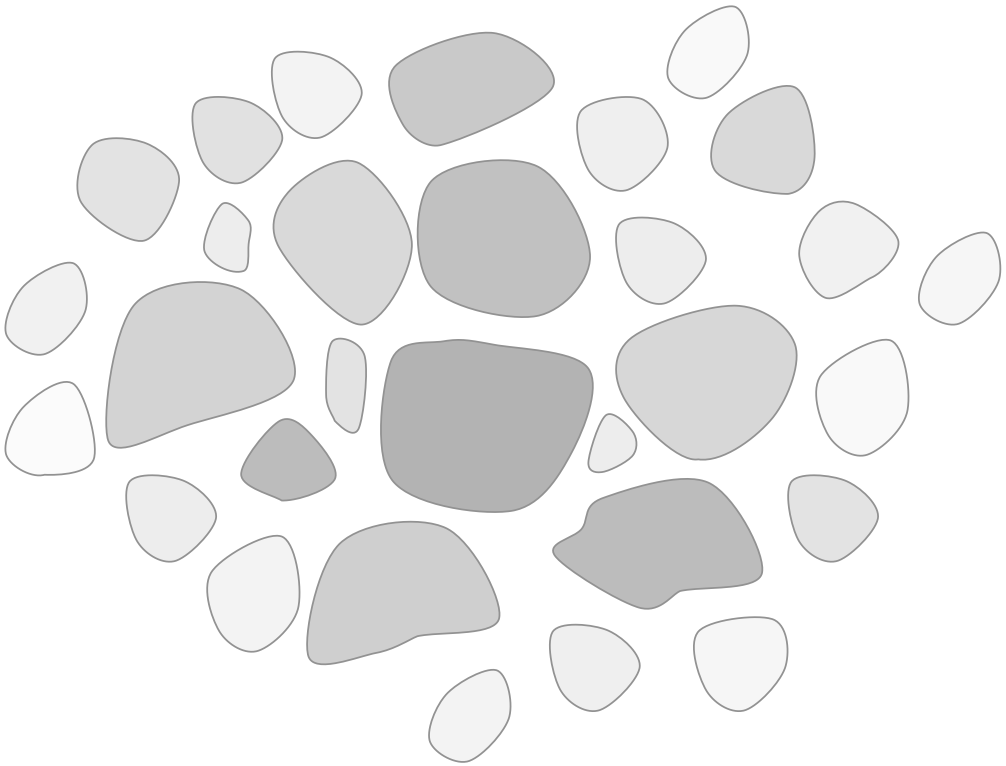
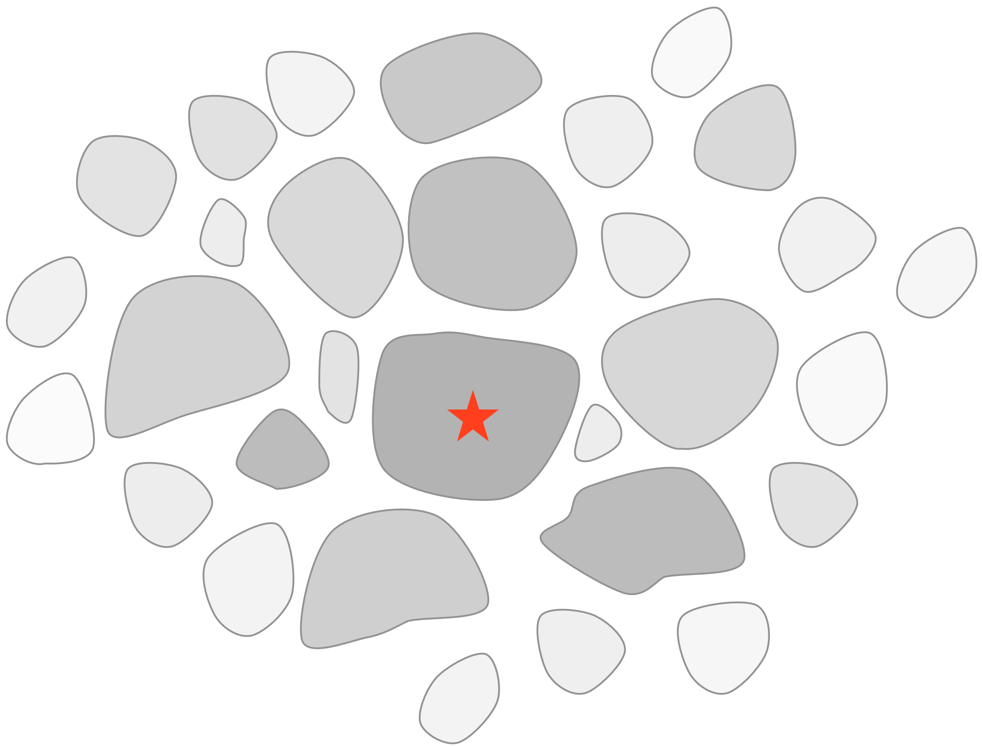
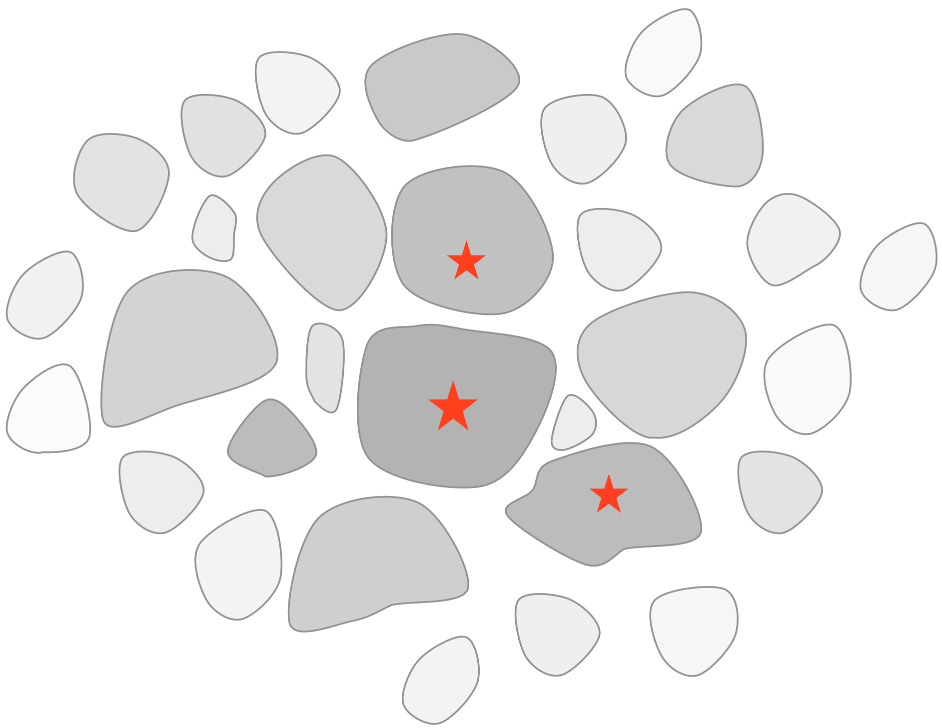


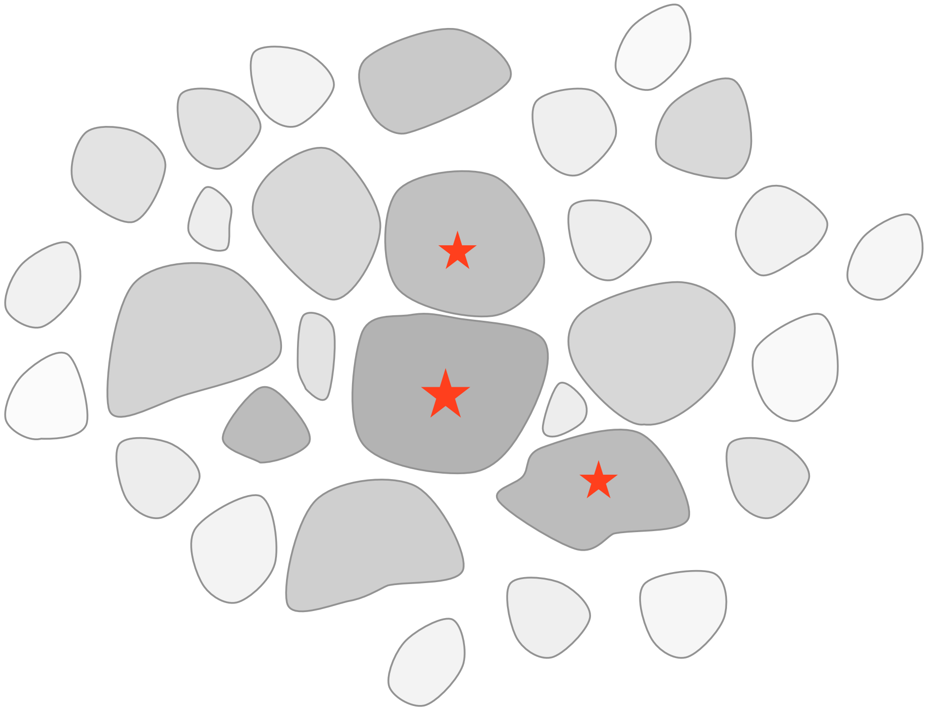
while region contracts, individual clumps collapse to form stars



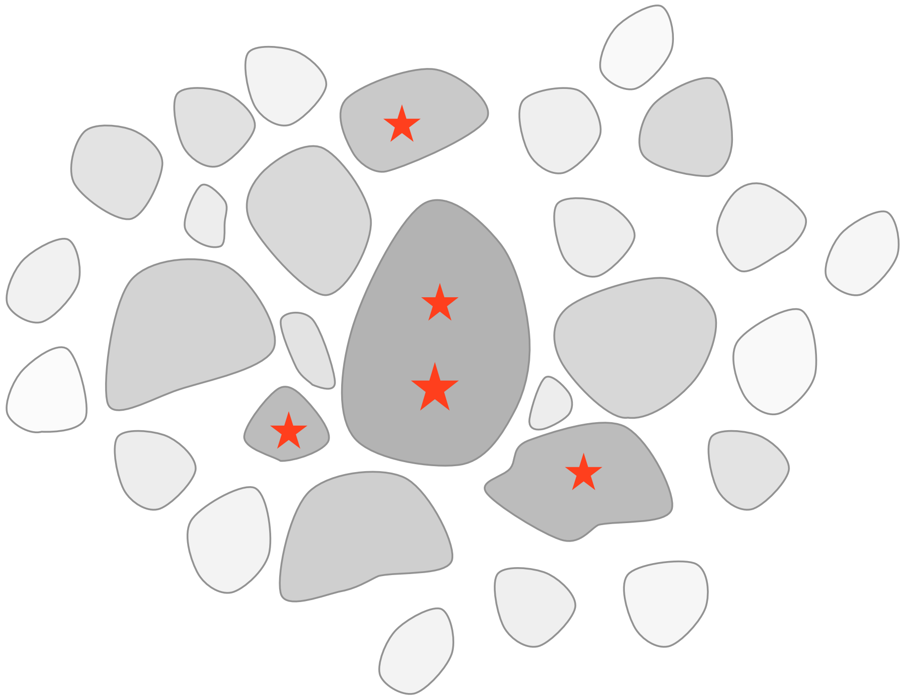
while region contracts, individual clumps collapse to form stars



individual clumps collapse to form stars

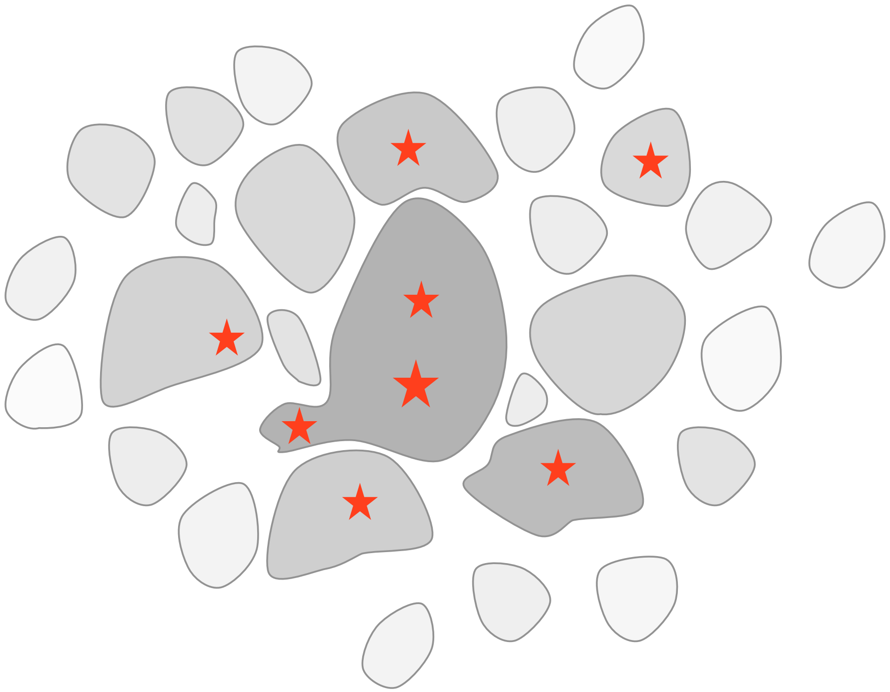


individual clumps collapse to form stars



$\alpha = E_{kin}/|E_{pot}| < 1$

in *dense clusters*, clumps may merge while collapsing
--> then contain multiple protostars



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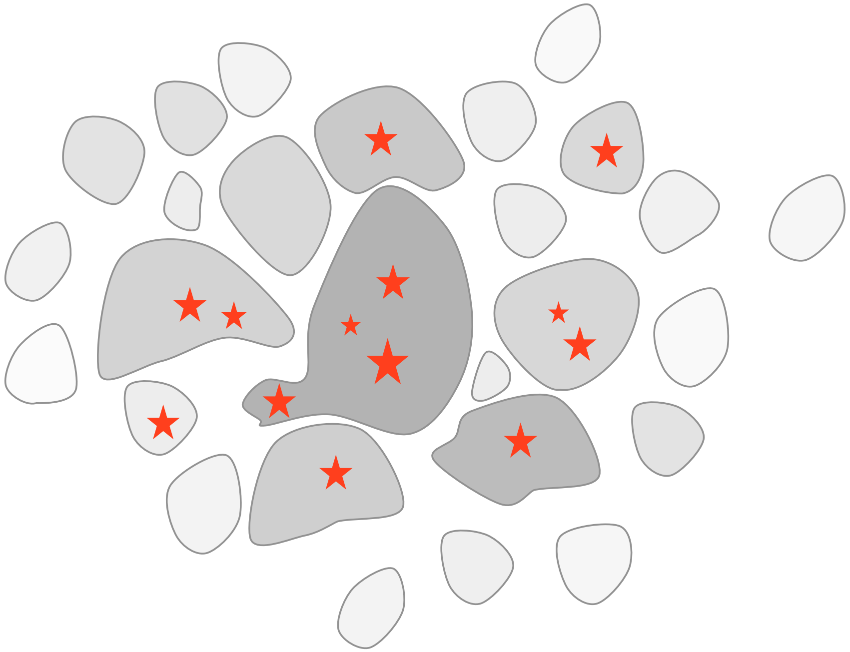
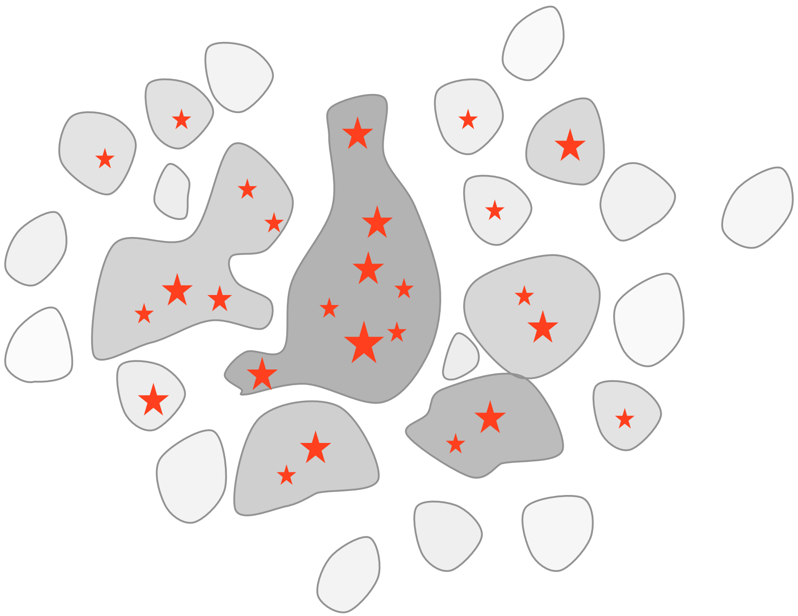
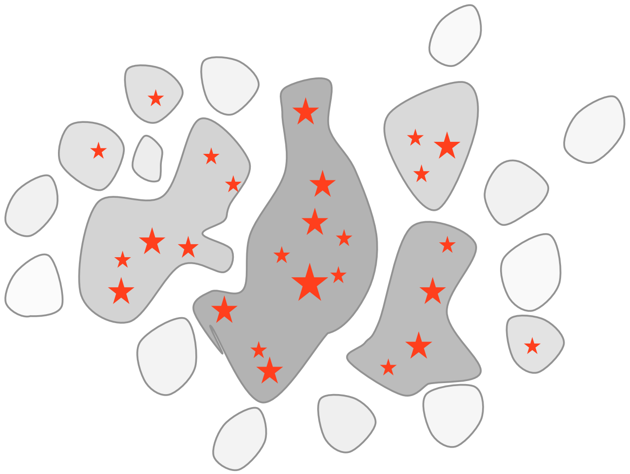


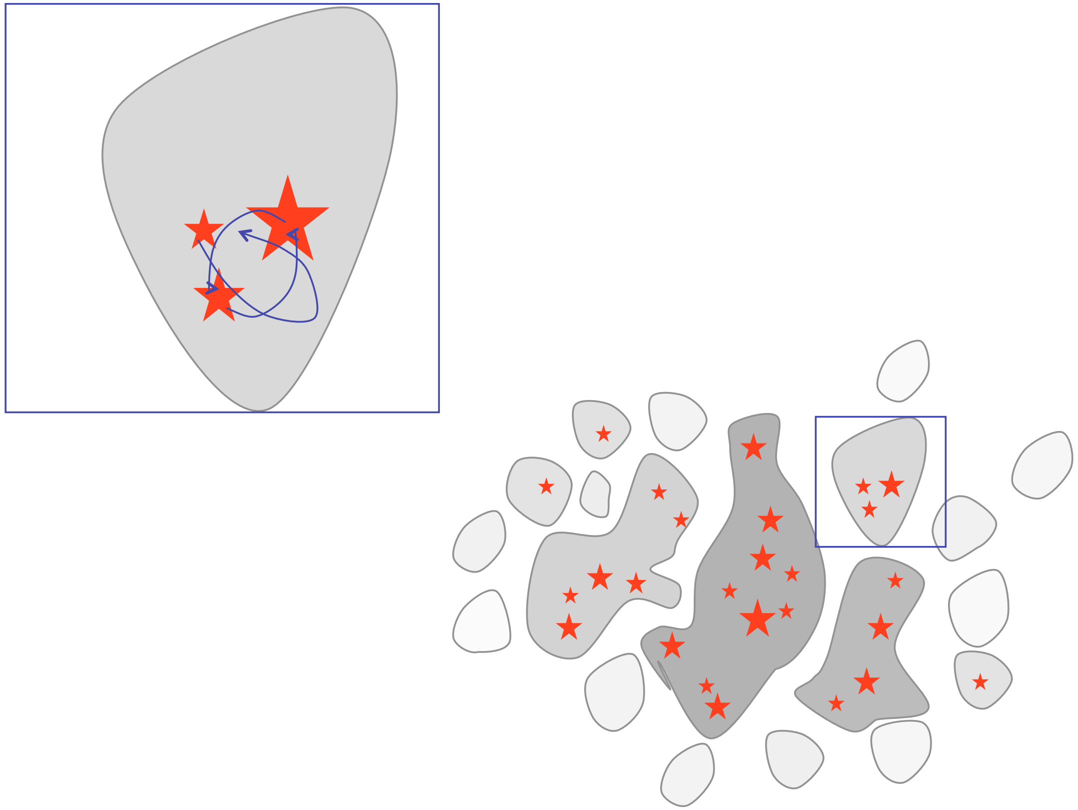
in *dense clusters*, clumps may merge while collapsing
--> then contain multiple protostars



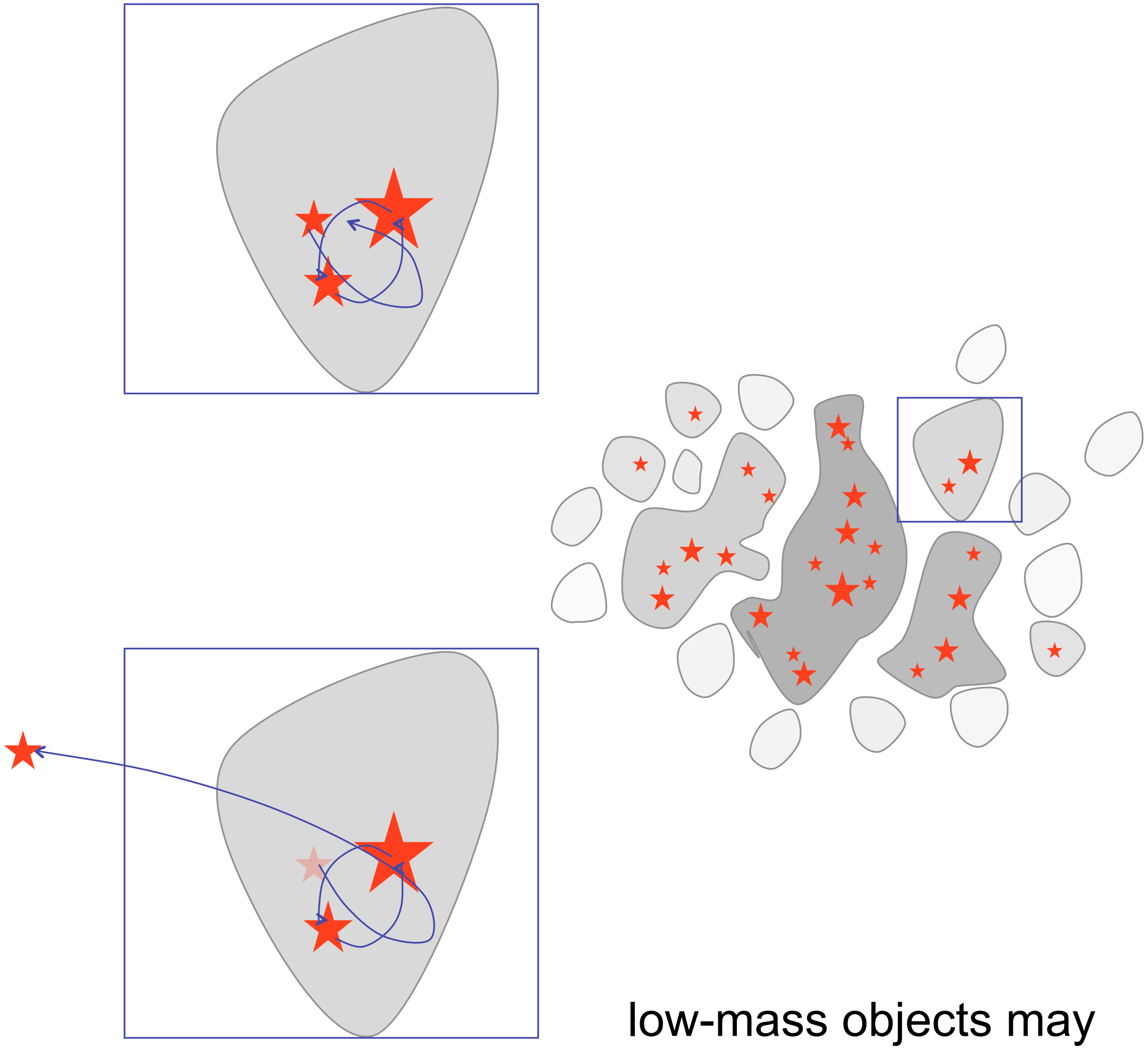
in *dense clusters*, competitive mass growth becomes important



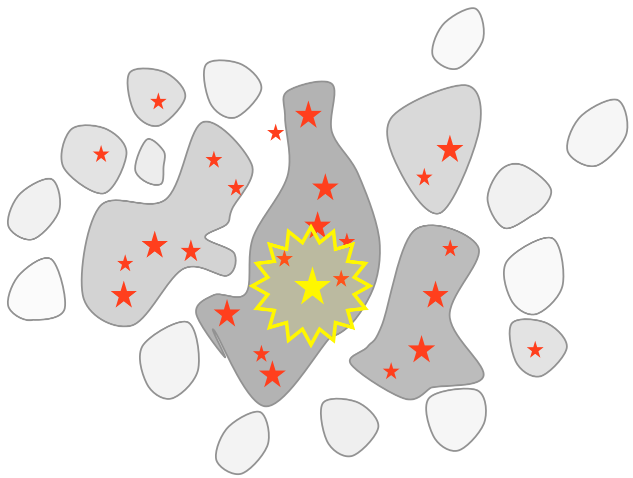
in *dense clusters*, competitive mass growth becomes important



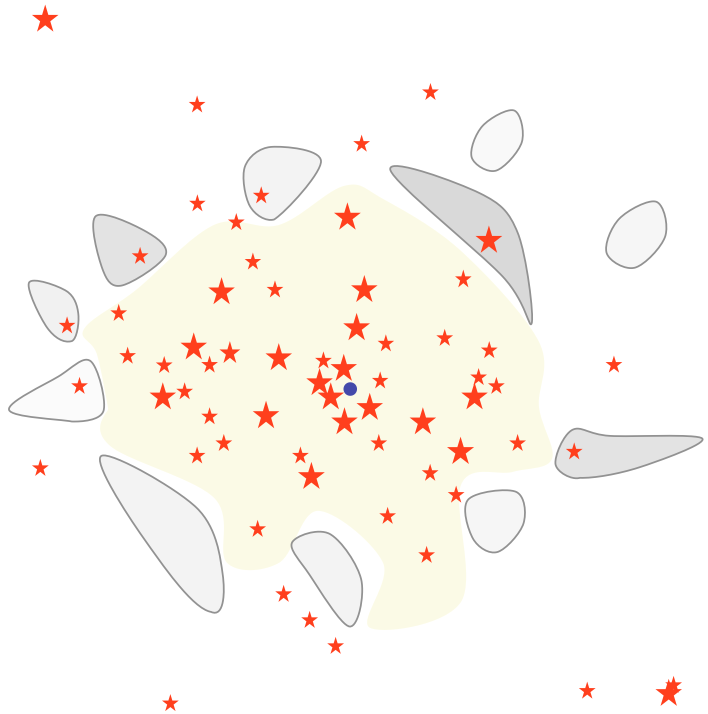
in *dense clusters*, *N*-body effects influence mass growth



low-mass objects may
become ejected --> accretion stops



feedback terminates star formation



result: *star cluster*, possibly with HII region


NGC 602 in the LMC: Hubble Heritage Image

result: *star cluster* with HII region

initial mass function

IMF

- distribution of stellar masses depends on
  - turbulent initial conditions
    --> mass spectrum of prestellar cloud cores
  - collapse and interaction of prestellar cores
    --> competitive accretion and *N*-body effects
  - thermodynamic properties of gas
    --> balance between heating and cooling
    --> EOS (determines which cores go into collapse)
  - (proto) stellar feedback terminates star formation
    ionizing radiation, bipolar outflows, winds, SN

(e.g. Larson 2003, Prog. Rep. Phys.; Mac Low & Klessen, 2004, Rev. Mod. Phys, 76, 125 - 194)

IMF

- distribution of stellar masses depends on
  - *turbulent initial conditions*
    --> mass spectrum of prestellar cloud cores ???
  - collapse and interaction of prestellar cores
    --> competitive accretion and *N*-body effects
  - thermodynamic properties of gas
    --> balance between heating and cooling
    --> EOS (determines which cores go into collapse)
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    ionizing radiation, bipolar outflows, winds, SN

(e.g. Larson 2003, Prog. Rep. Phys.; Mac Low & Klessen, 2004, Rev. Mod. Phys, 76, 125 - 194)

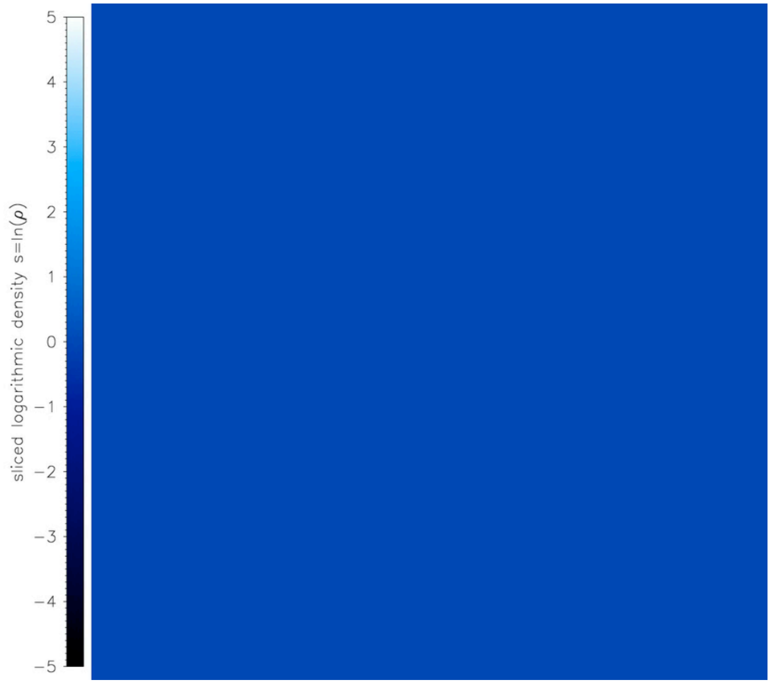
compressive vs. rotational driving

- statistical characteristics of turbulence depend strongly on „type“ of driving
- example: dilatational vs. solenoidal driving
- question: what drives ISM turbulence on different scales?

dilatational vs. solenoidal

density as function of time / cut through 1024^3 cube simulation (FLASH)

compressive
larger structures, higher ρ-contrast



rotational
smaller structures, small ρ-pdf

Federrath, Klessen, Schmidt (2008a,b)

dilatational vs. solenoidal

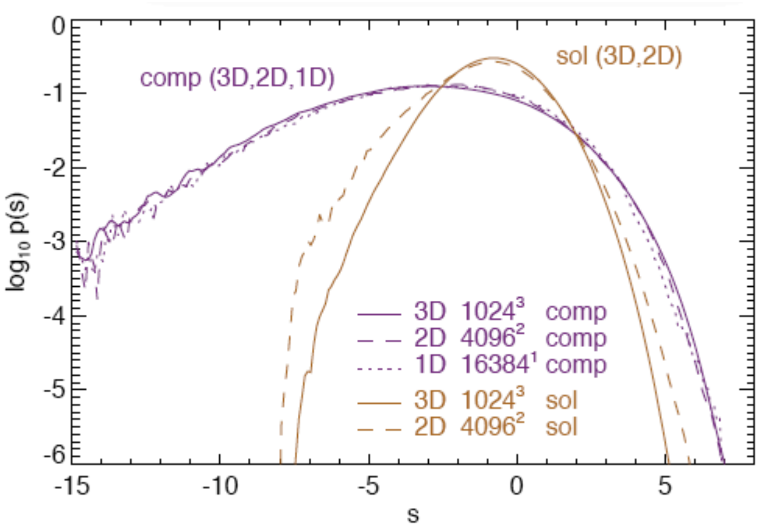


FIG. 3.— Volume-weighted density PDFs $p(s)$ obtained from 3D, 2D and 1D simulations with compressive forcing and from 3D and 2D simulations using solenoidal forcing. Note that in 1D, only compressive forcing is possible as in the study by Passot & Vázquez-Semadeni (1998). As suggested by eq. (5), compressive forcing yields almost identical density PDFs in 1D, 2D and 3D with $b \sim 1$, whereas solenoidal forcing leads to a density PDF with $b \sim 1/2$ in 2D and with $b \sim 1/3$ in 3D.

Federrath, Klessen, Schmidt (2008a)

- density pdf depends on "dimensionality" of driving
  - relation between width of pdf and Mach number

$$\sigma_\rho / \rho_0 = b\mathcal{M}$$

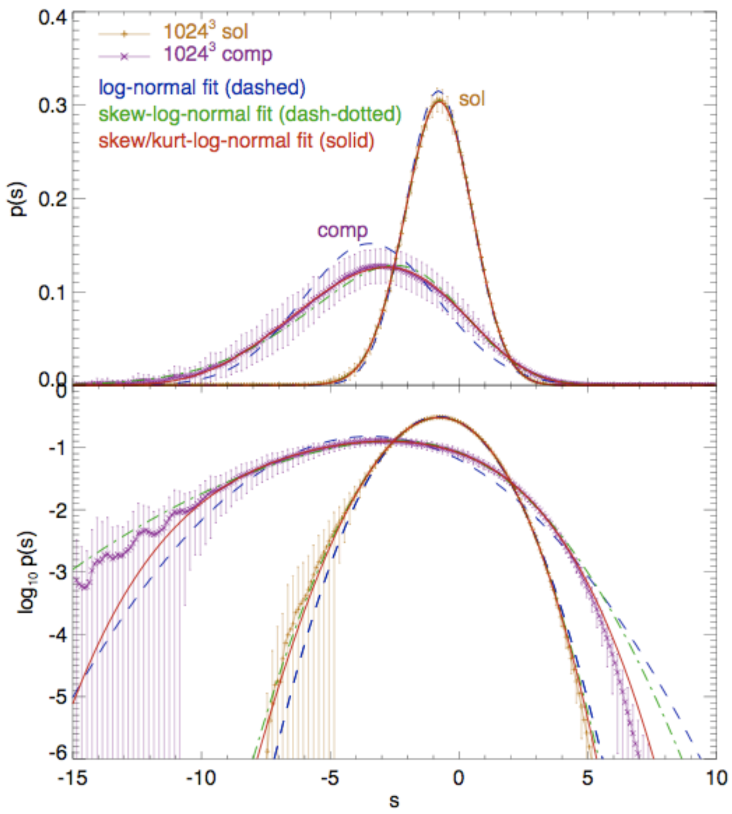
  - with b depending on ζ via

$$b = 1 + \left[\frac{1}{D} - 1\right]\zeta = \begin{cases} 1 - \frac{2}{3}\zeta & \text{, for } D = 3 \\ 1 - \frac{1}{2}\zeta & \text{, for } D = 2 \\ 1 & \text{, for } D = 1 \end{cases}$$

  - with ζ being the ratio of dilatational vs. solenoidal modes:

$$\mathcal{P}_{ij}^{\zeta} = \zeta \mathcal{P}_{ij}^{\perp} + (1-\zeta)\mathcal{P}_{ij}^{\|} = \zeta\delta_{ij} + (1-2\zeta)\frac{k_i k_j}{|k|^2}$$

dilatational vs. solenoidal

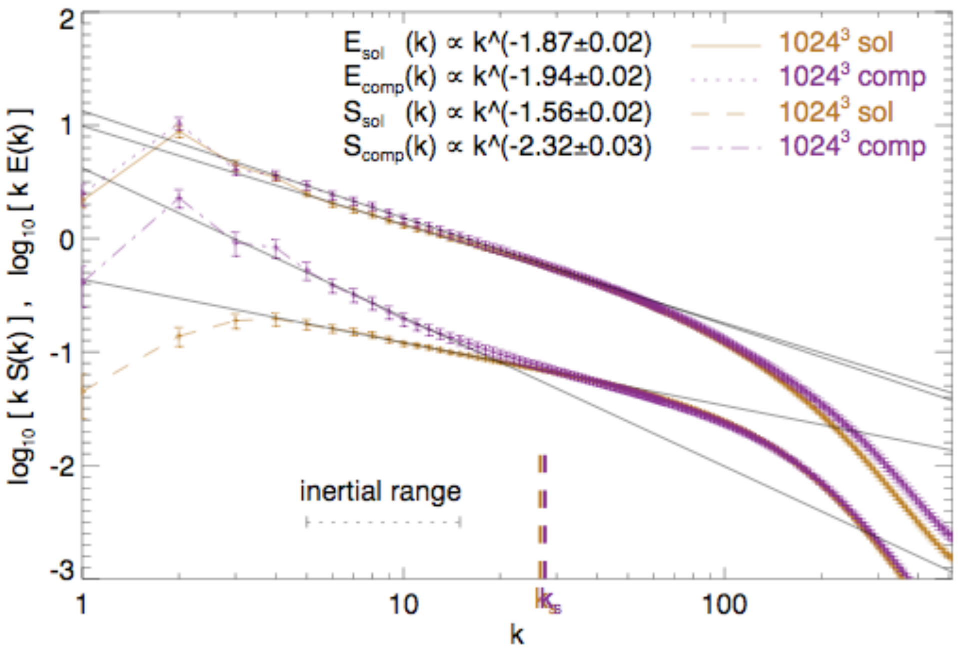


good fit needs 3rd and 4th moment of distribution!

Federrath, Klessen, Schmidt (2008b)

- density pdf depends on "dimensionality" of driving
  → is that a problem for the Krumholz & McKee model of the SF efficiency?
- density pdf of compressive driving is *NOT log-normal*
  → is that a problem for the Padoan & Nordlund, or Hennebelle & Chabrier IMF model?
- most "physical" sources should be *compressive* (convergent flows from spiral shocks or SN)

dilatational vs. solenoidal



compensated density spectrum kS(k) shows clear break at sonic scale. below that shock compression no longer is important in shaping the power spectrum ...

- density power spectrum differs between dilatational and solenoidal driving!

→ dilatational driving leads to break at sonic scale!

- can we use that to determine driving sources from observations ?

Federrath, Klessen, Schmidt (2008b)

IMF

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    ionizing radiation, bipolar outflows, winds, SN

(e.g. Larson 2003, Prog. Rep. Phys.; Mac Low & Klessen, 2004, Rev. Mod. Phys, 76, 125 - 194)

example: model of Orion cloud

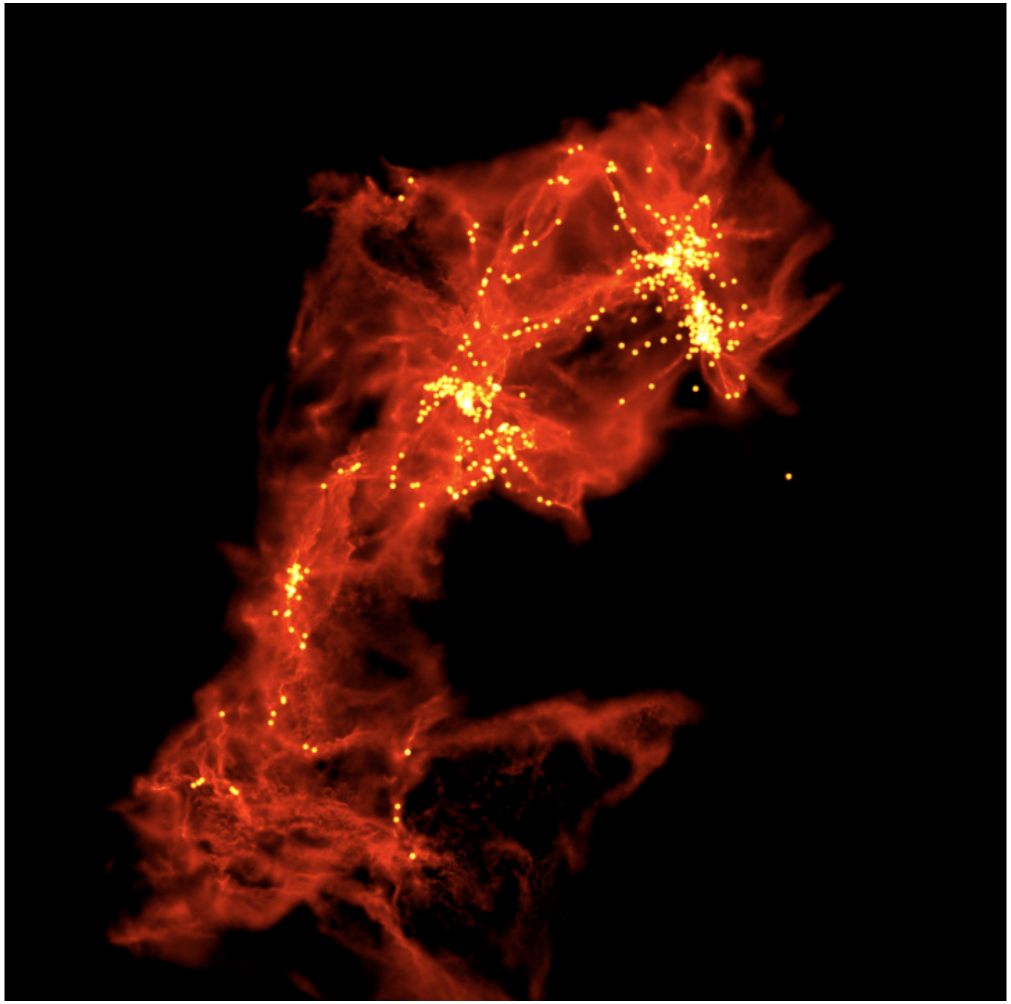
„model“ of Orion cloud:
15.000.000 SPH particles,
10^4 M_{sun} in 10 pc, mass resolution
0,02 M_{sun}, forms ~2.500
„stars“ (sink particles)

isothermal EOS, top bound, bottom unbound

has clustered as well as distributed „star“ formation

efficiency varies from 1% to 20%

develops full IMF
(distribution of sink particle masses)



(Bonnell & Clark 2008)

example: model of Orion cloud

„model" of Orion cloud:
15.000.000 SPH particles,
10^4 M_{sun} in 10 pc, mass resolution
0,02 M_{sun}, forms ~2.500
„stars" (sink particles)

MASSIVE STARS

- form early in high-density gas clumps (cluster center)
- high accretion rates, maintained for a long time

LOW-MASS STARS

- form later as gas falls into potential well
- high relative velocities
- little subsequent accretion

Bonnell & Clark 2008

example: model of Orion cloud

„model“ of Orion cloud:
15.000.000 SPH particles,
10^4 M_{sun} in 10 pc, mass resolution
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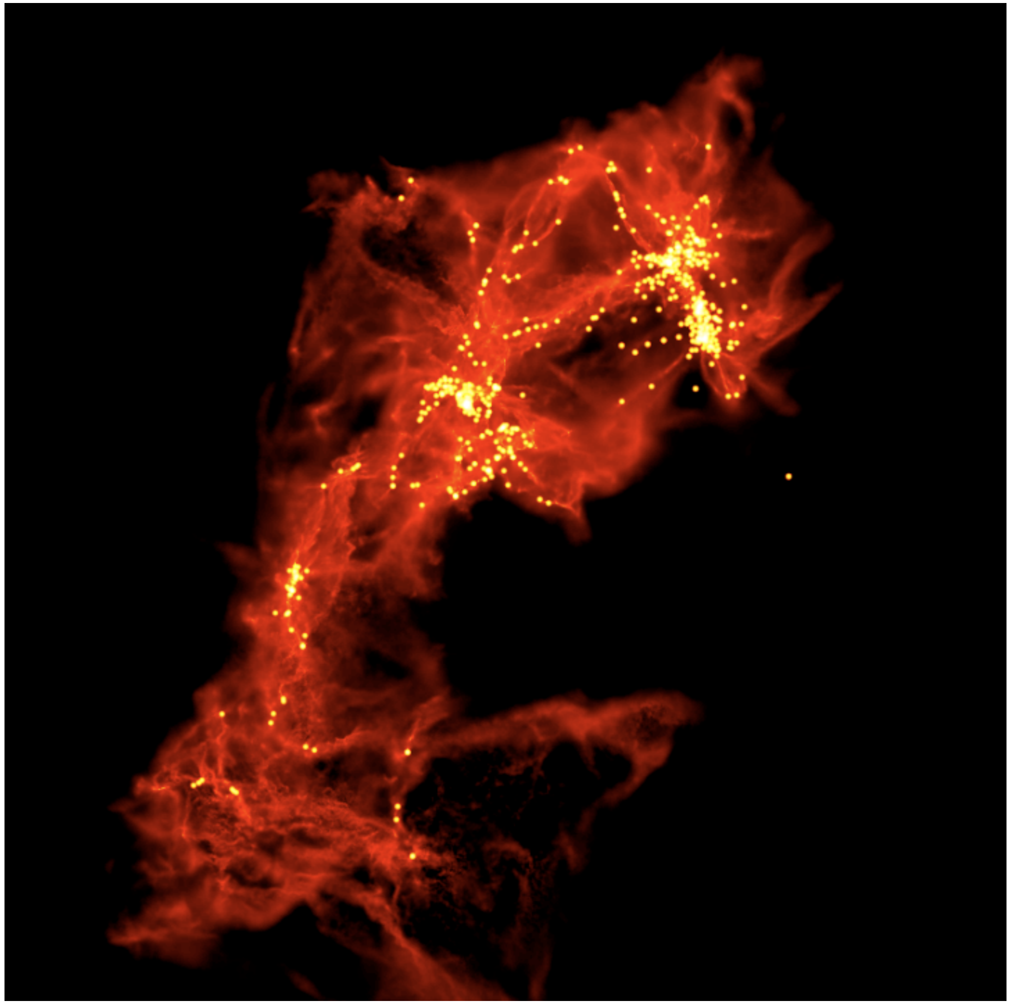
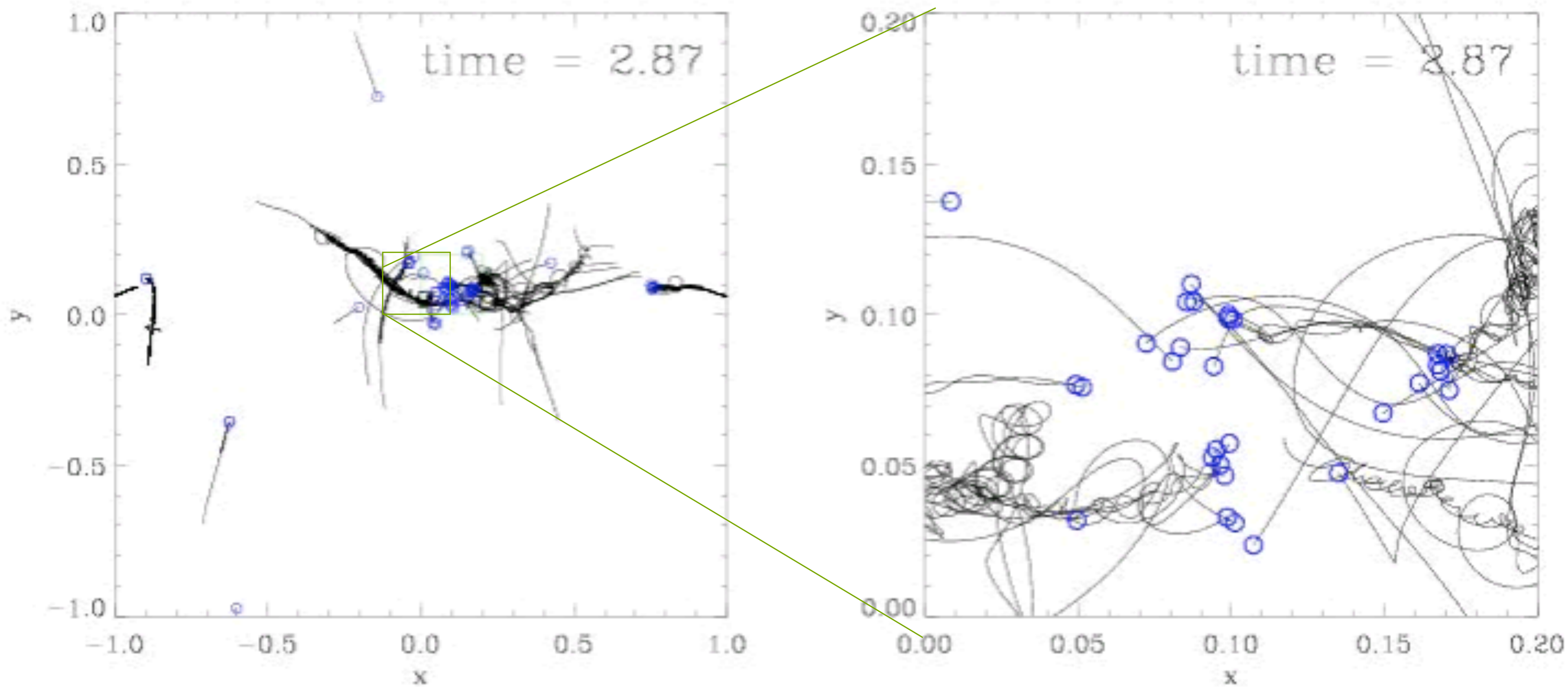


(Bonnell & Clark 2008)

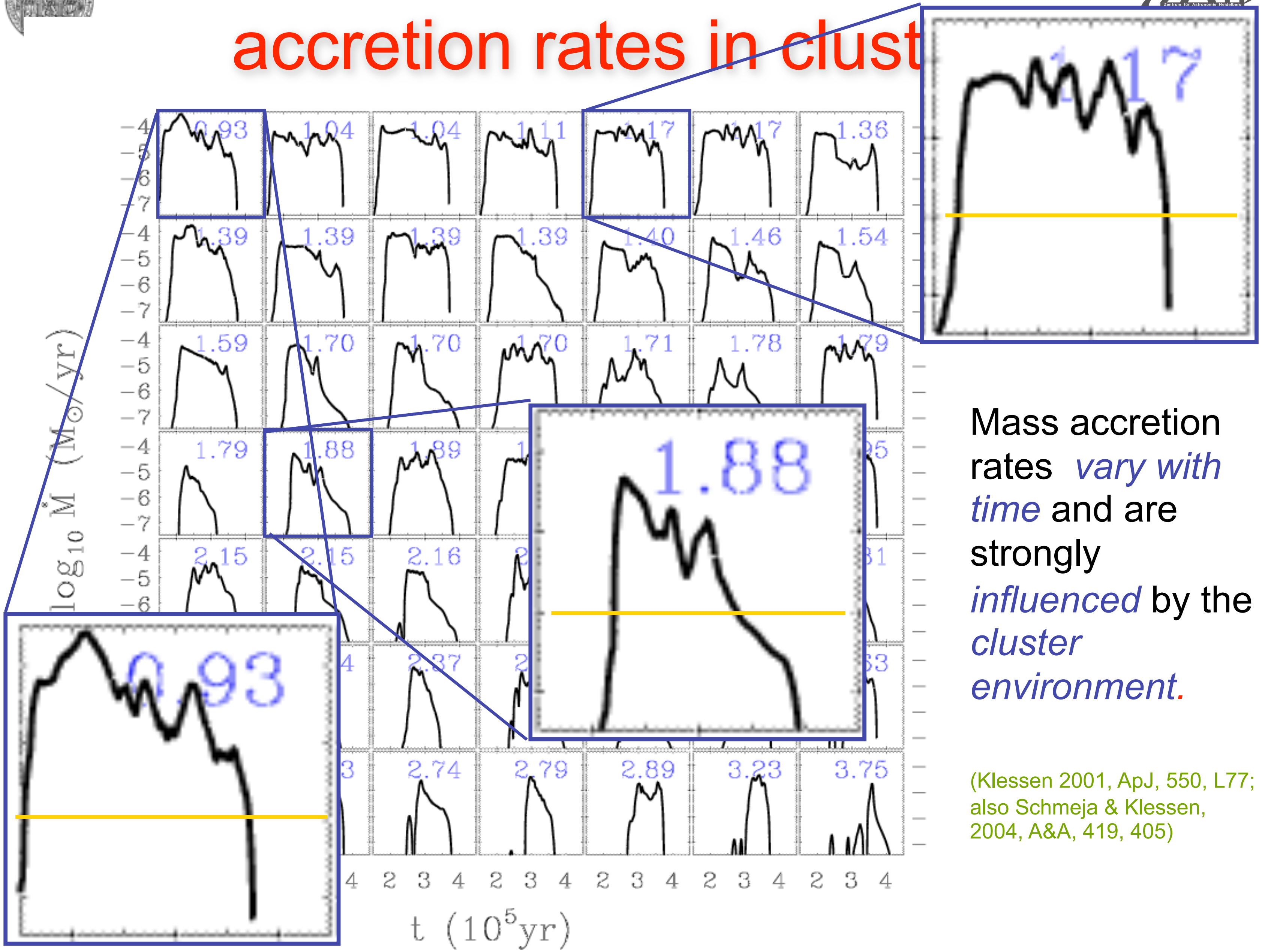
Dynamics of nascent star cluster

in dense clusters protostellar interaction may be come important!



Trajectories of protostars in a nascent dense cluster created by gravoturbulent fragmentation
(from Klessen & Burkert 2000, ApJS, 128, 287)

accretion rates in clust



Mass accretion rates *vary with time* and are strongly *influenced* by the *cluster environment*.

(Klessen 2001, ApJ, 550, L77; also Schmeja & Klessen, 2004, A&A, 419, 405)

IMF

- distribution of stellar masses depends on
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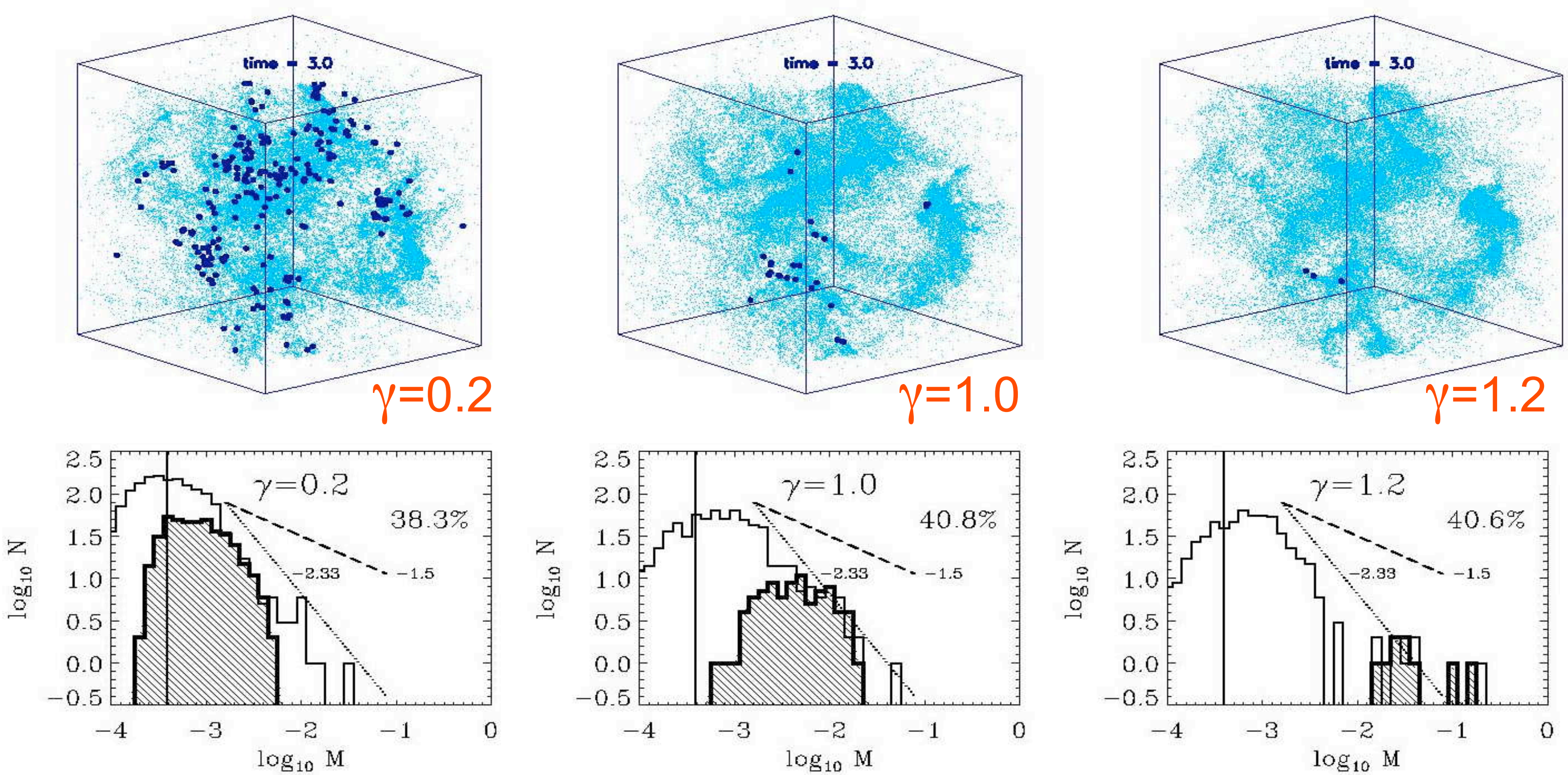
(e.g. Larson 2003, Prog. Rep. Phys.; Mac Low & Klessen, 2004, Rev. Mod. Phys, 76, 125 - 194)

dependency on EOS

- degree of fragmentation depends on *EOS!*

- polytropic EOS: **p** $\propto \rho^\gamma$
- $\gamma<1$: dense cluster of low-mass stars
- $\gamma>1$: isolated high-mass stars
- (see Li, Klessen, & Mac Low 2003, ApJ, 592, 975; also Kawachi & Hanawa 1998, Larson 2003)

dependency on EOS



for γ<1 fragmentation is enhanced → *cluster of low-mass stars*

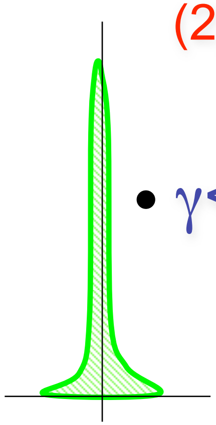
for γ>1 it is suppressed → formation of *isolated massive stars*

(from Li, Klessen, & Mac Low 2003, ApJ, 592, 975)

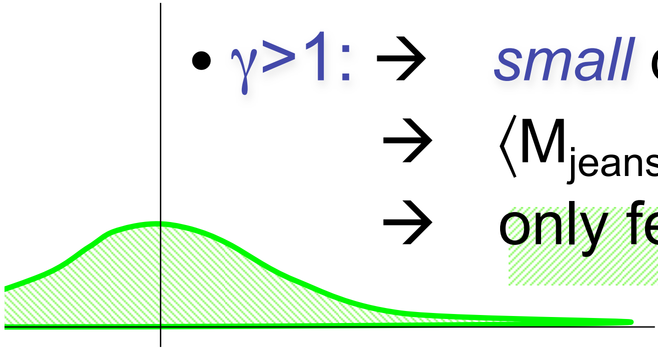
how does that work?

(1) $\mathbf{p} \propto \rho^{\gamma} \rightarrow \rho \propto \mathbf{p}^{1/\gamma}$

(2) $\mathbf{M_{jeans}} \propto \gamma^{3/2}\, \rho^{(3\gamma-4)/2}$



- $\gamma<1$: → *large* density excursion for given pressure
  → $\langle M_{jeans} \rangle$ becomes small
  → number of fluctuations with $M > M_{jeans}$ is large

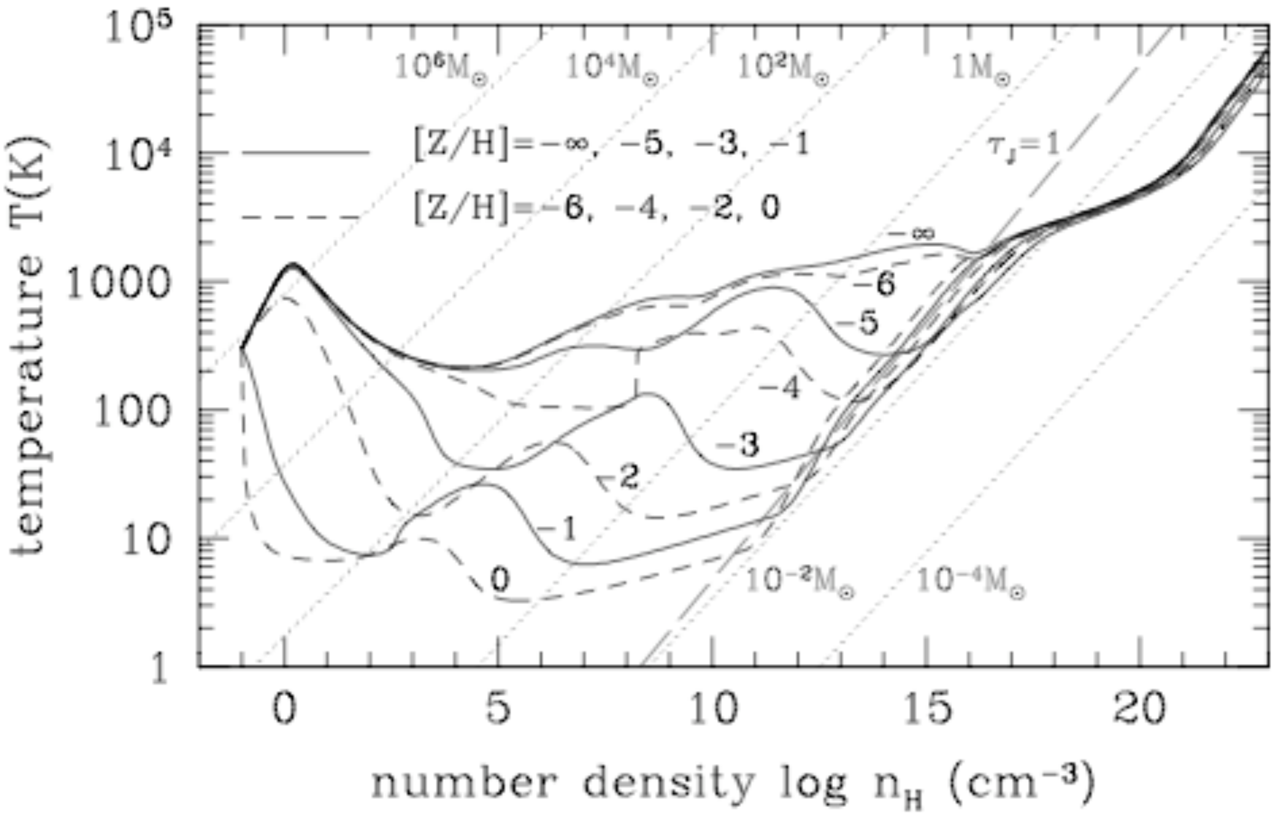


- $\gamma>1$: → *small* density excursion for given pressure
  → $\langle M_{jeans} \rangle$ is large
  → only few and massive clumps exceed M_{jeans}

EOS in different environments

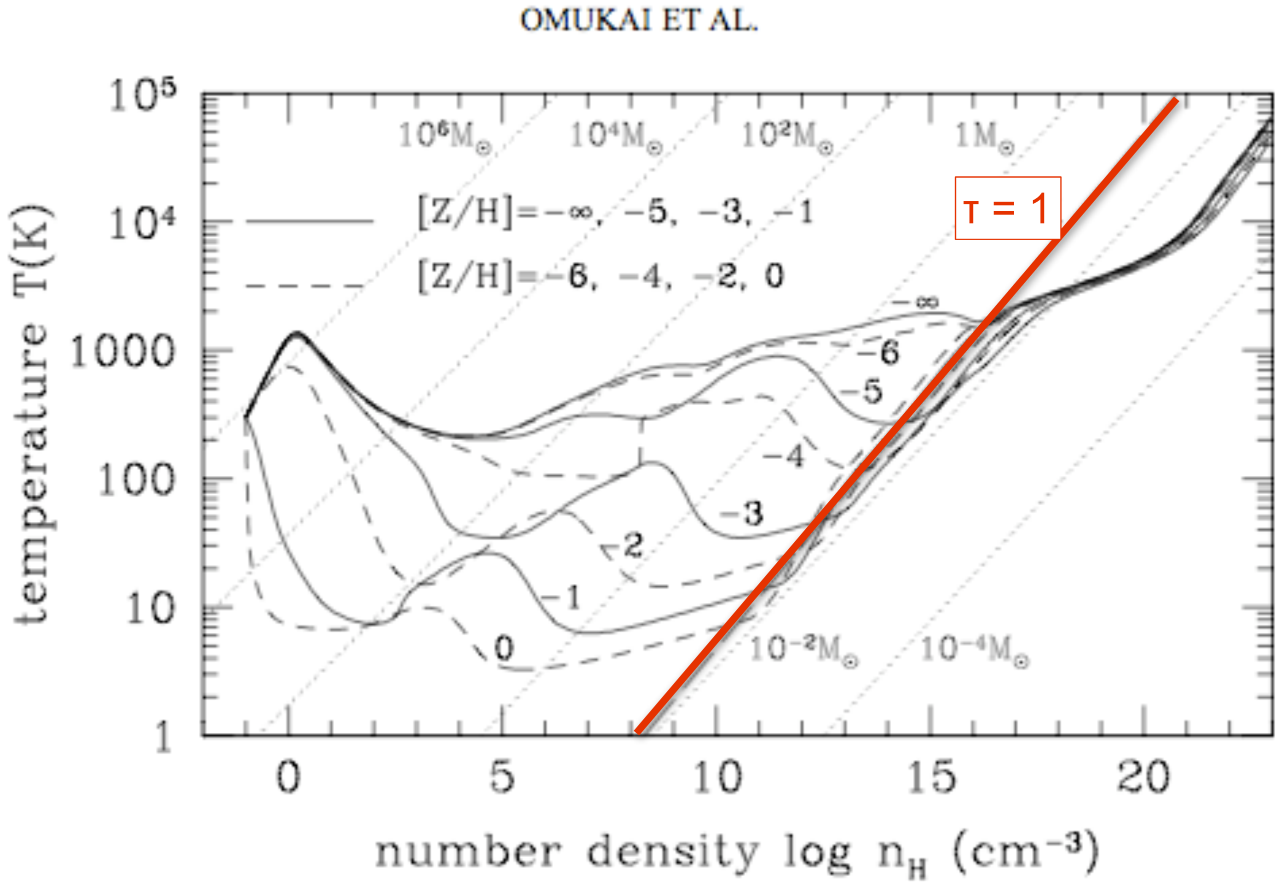
EOS as function of metallicity

OMUKAI ET AL.



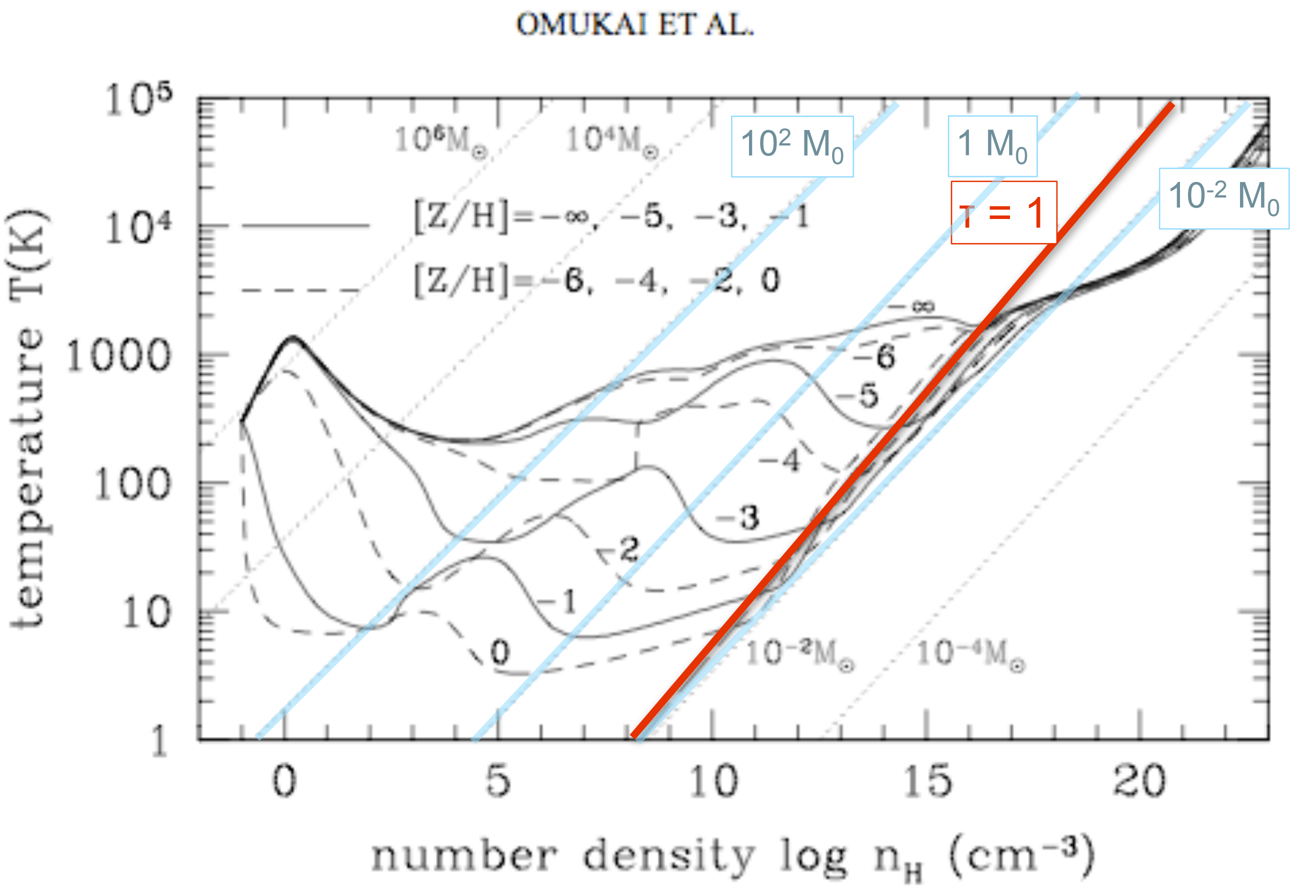
(Omukai et al. 2005)

EOS as function of metallicity



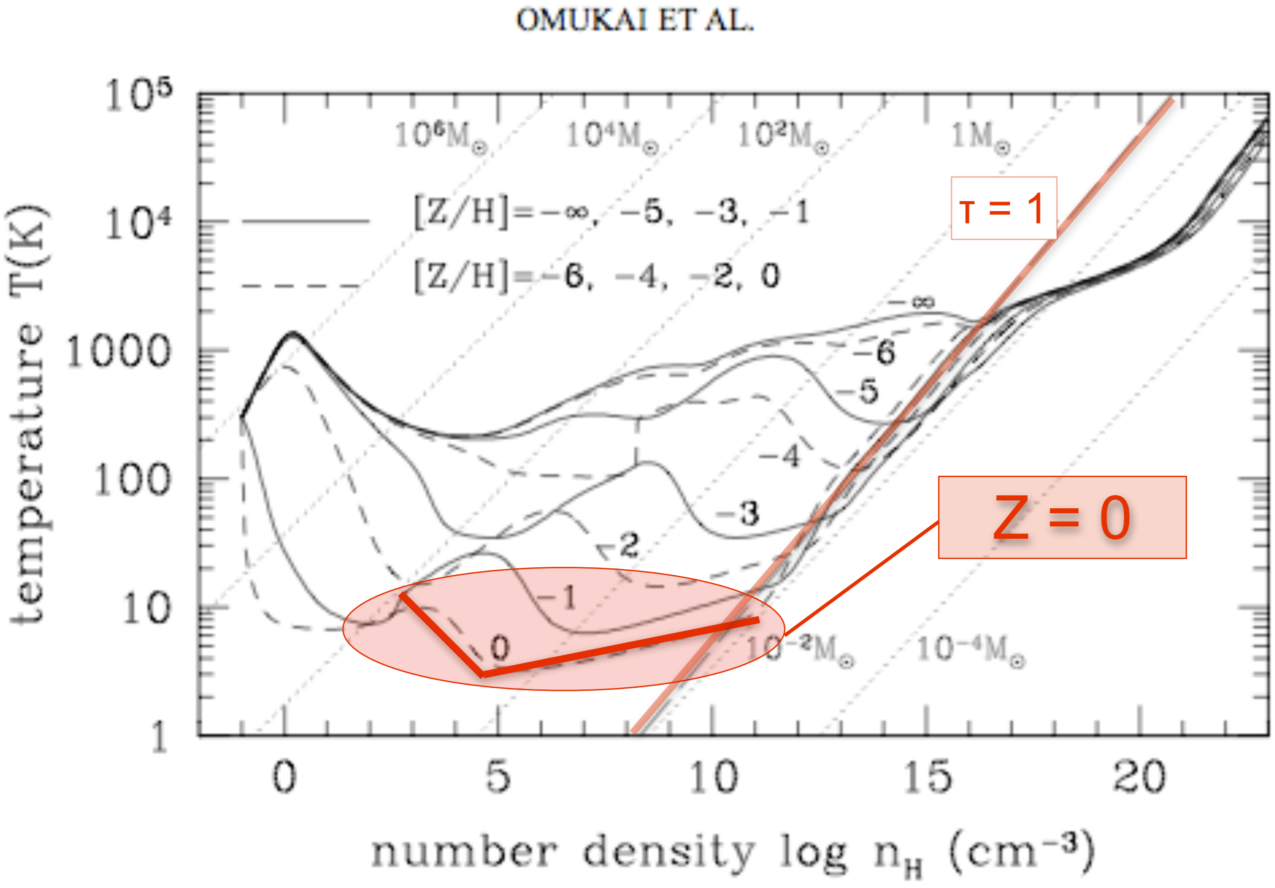
(Omukai et al. 2005)

EOS as function of metallicity



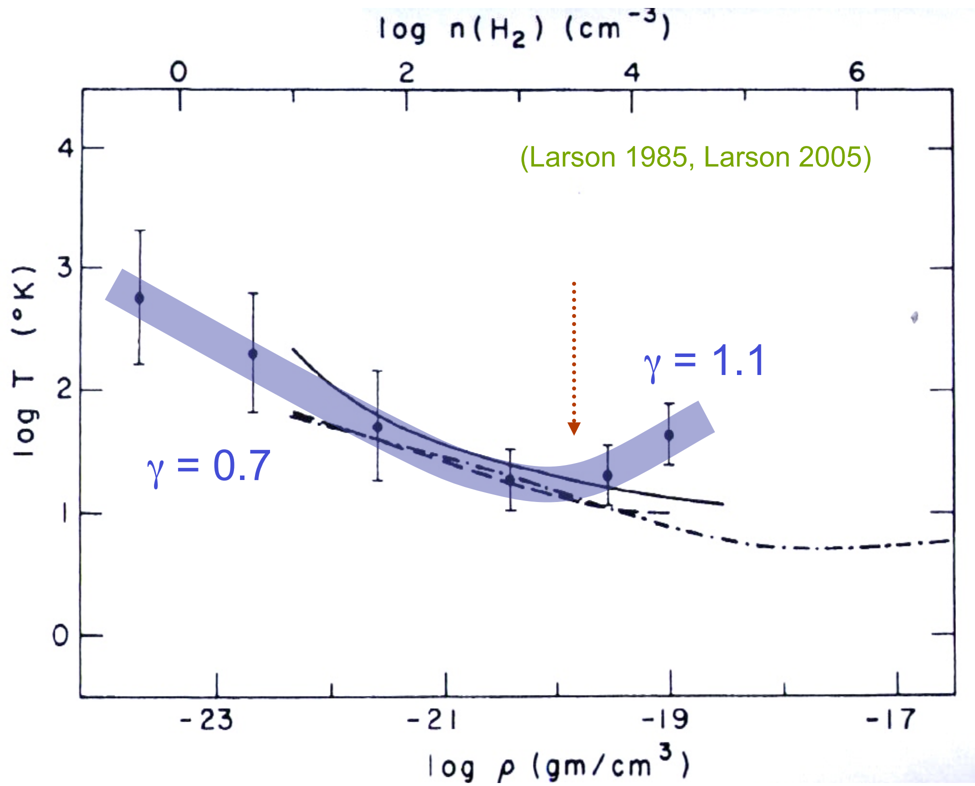
(Omukai et al. 2005)

present-day star formation



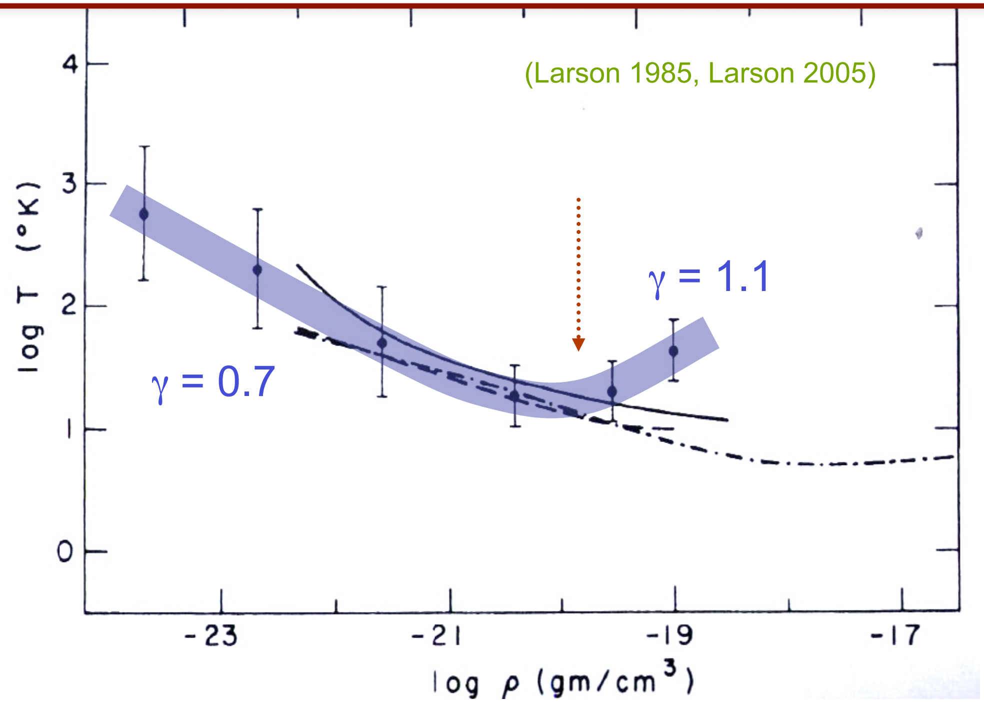
(Omukai et al. 2005, Jappsen et al. 2005, Larson 2005)

present-day star formation



present-day star formation

This kink in EOS is very insensitive to environmental conditions such as ambient radiation field
--> reason for universal for of the IMF? (Elmegreen et al. 2008)

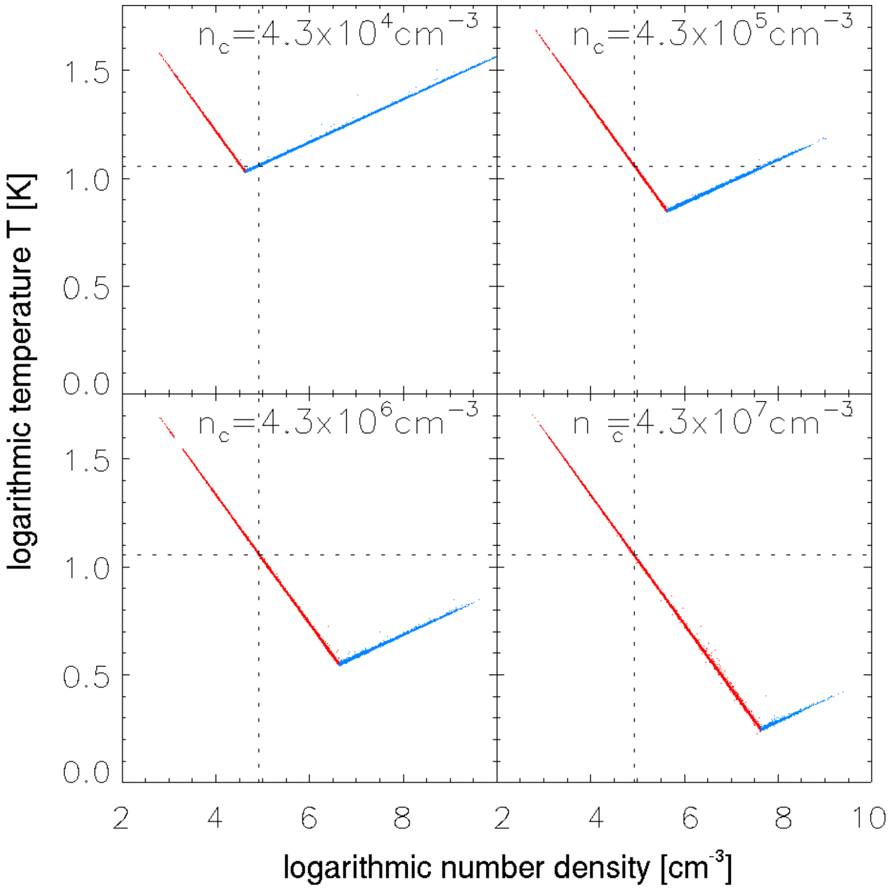


IMF from simple piece-wise polytropic EOS

$\gamma_1 = 0.7$

$\gamma_2 = 1.1$

$T \sim \rho^{\gamma-1}$



EOS and Jeans Mass:

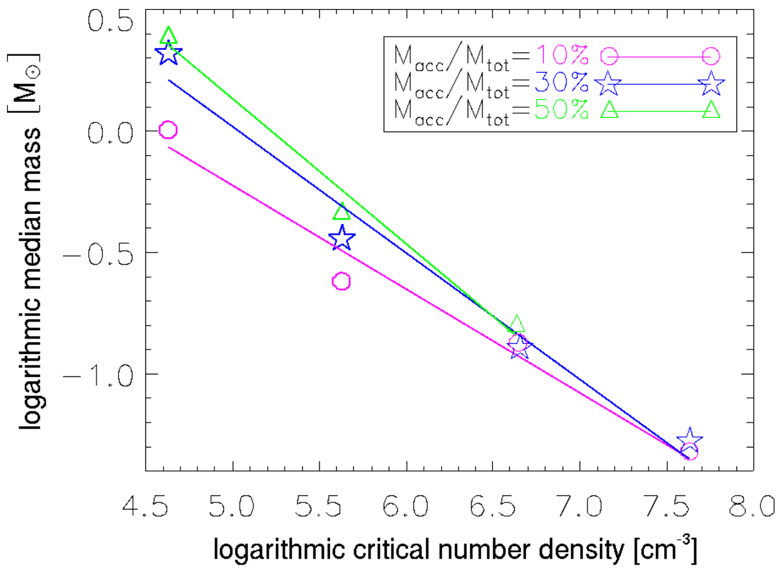
$p \propto \rho^\gamma \rightarrow \rho \propto p^{1/\gamma}$

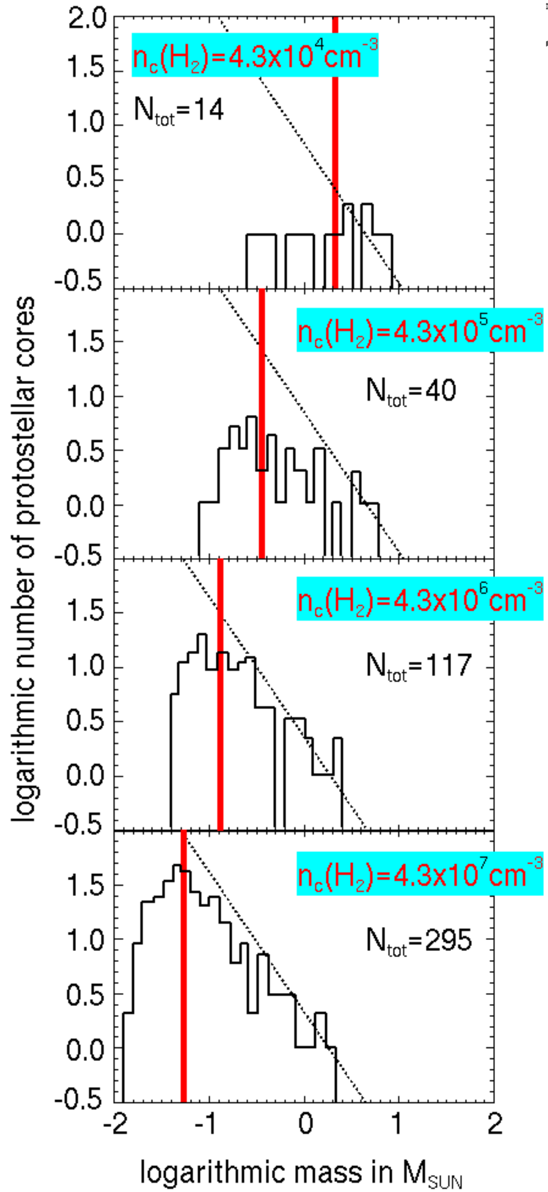
$M_{jeans} \propto \gamma^{3/2} \rho^{(3\gamma-4)/2}$

(Jappsen et al. 2005)

IMF from simple piece-wise EOS

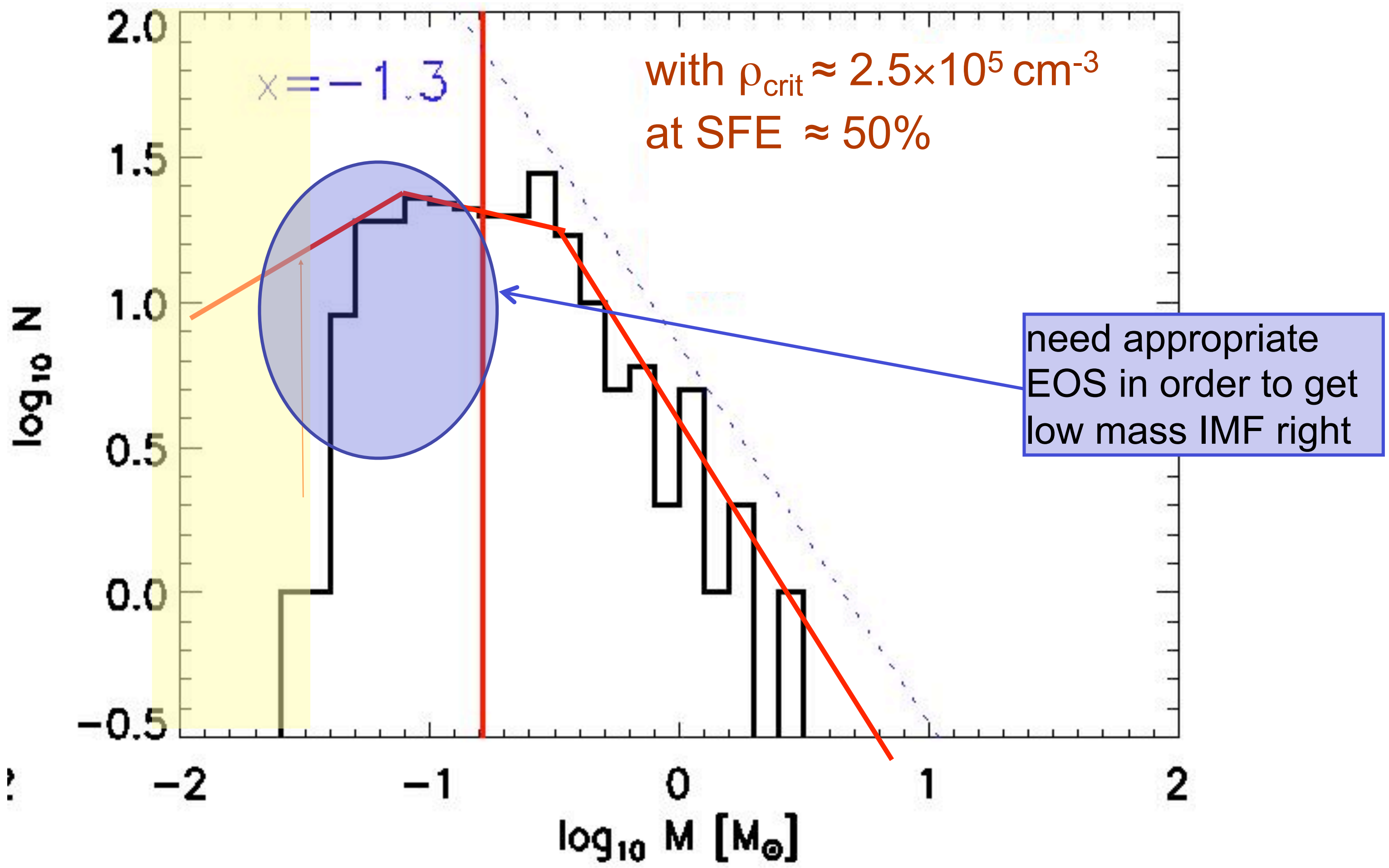
critical density ⇧ ⇨ median mass ⇩





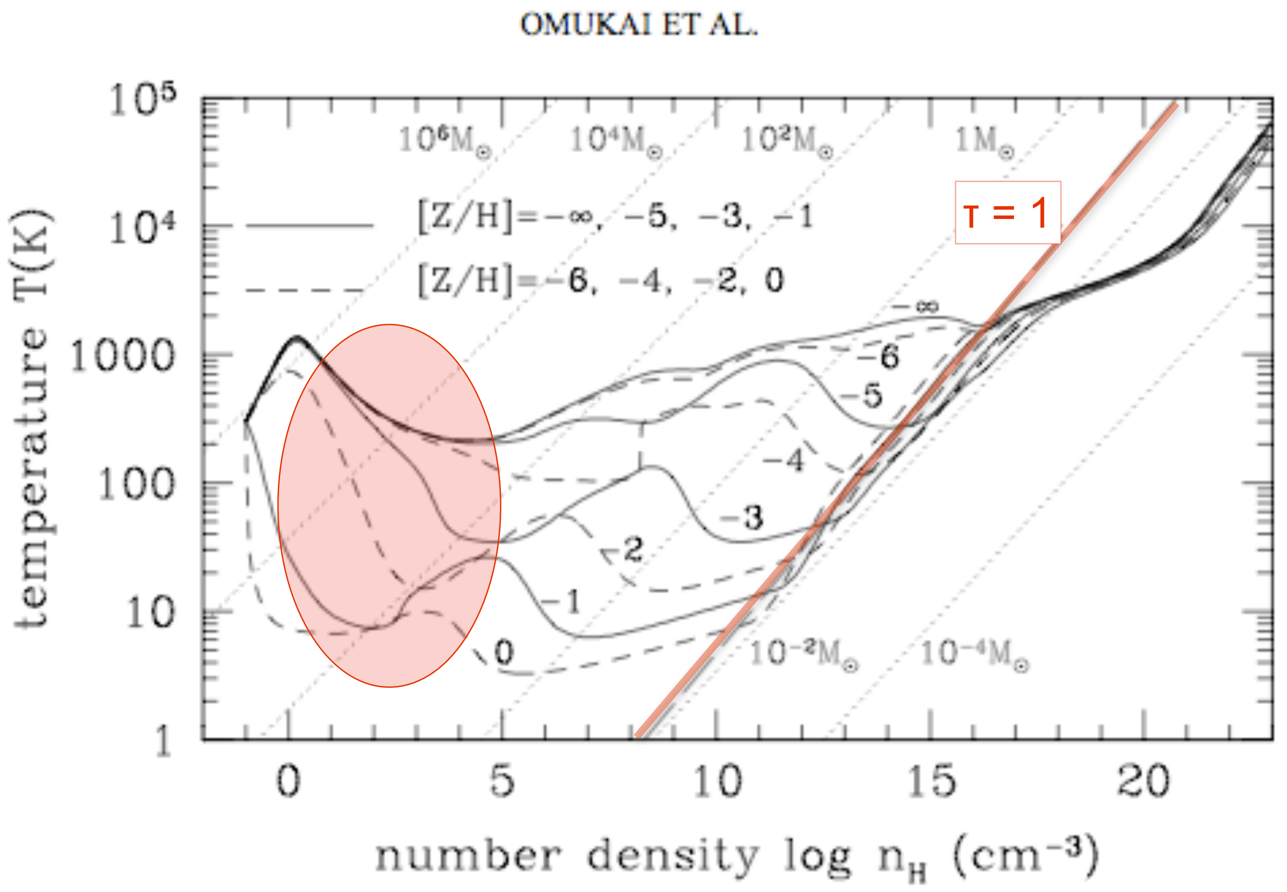
(Jappsen et al. 2005)

IMF in nearby molecular clouds



(Jappsen et al. 2005, A&A, 435, 611)

dependence on Z at low density



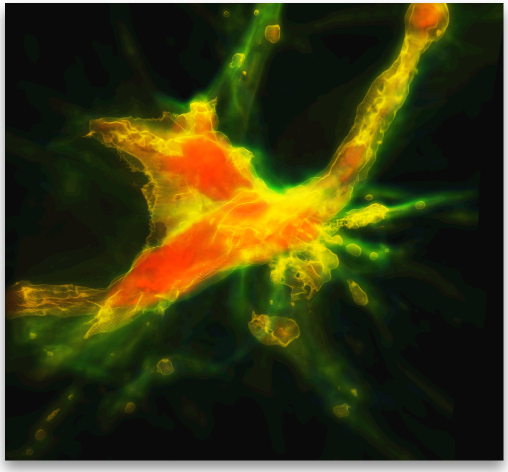
(Omukai et al. 2005)

dependence on Z at low density

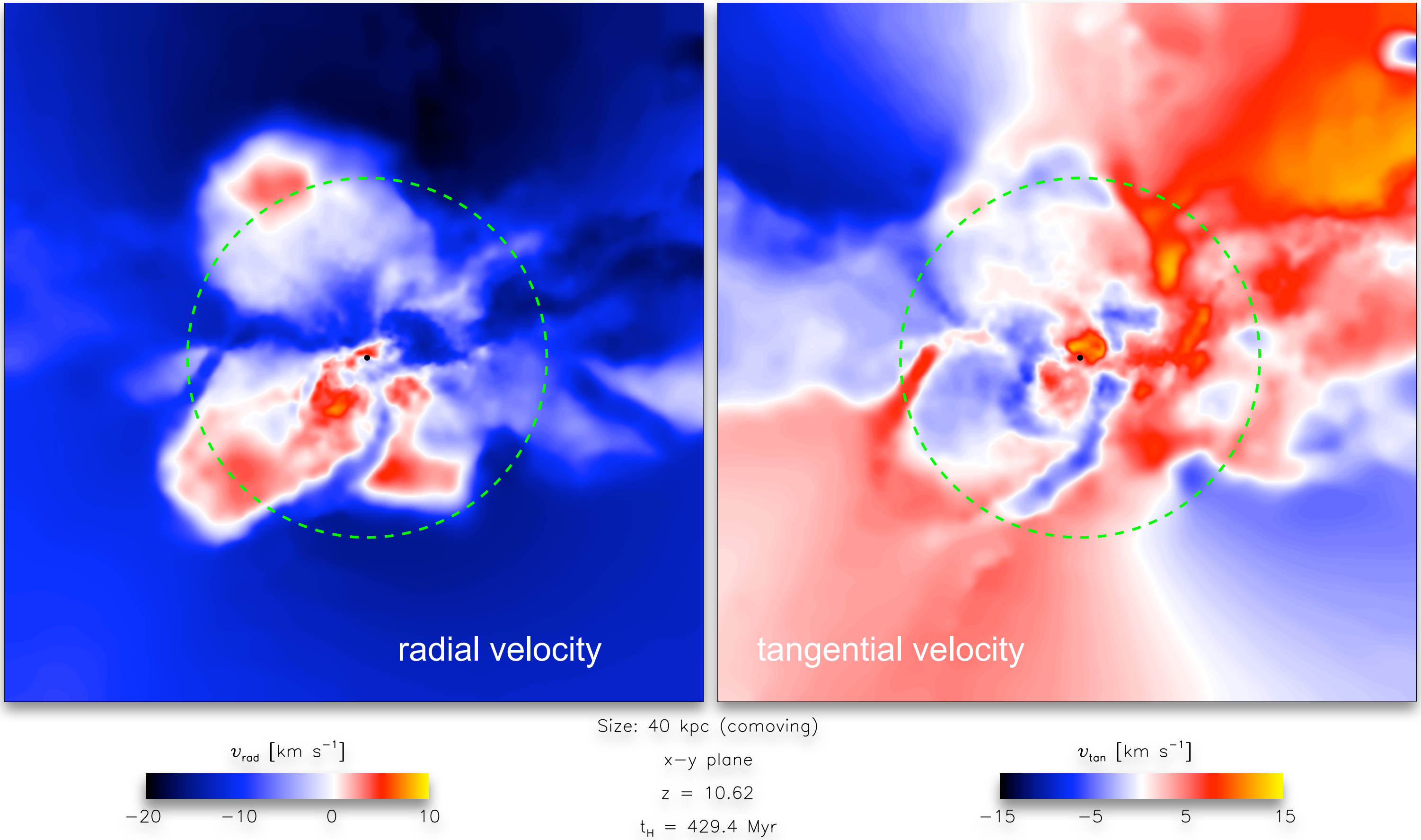
- at densities $n < 10^2$ cm^{-3} and metallicities $Z < 10^{-2}$ *H_2 cooling* dominates behavior.
  (Jappsen et al. 2007)
- fragmentation depends on *initial conditions*
  - example 1: *solid-body rotating top-hat* initial conditions with dark matter fluctuations (a la Bromm et al. 1999) fragment no matter what metallicity you take (in regime $n \leq 10^6$ cm^{-3}) because *unstable disk* builds up
    (Jappsen et al. 2009a)
  - example 2: *centrally concentrated halo* does *not* fragment up to densities of $n \approx 10^6$ cm^{-3} up to metallicities $Z \approx -1$ (Jappsen et al. 2009b)

implications for Pop III

- star formation will depend on *degree of turbulence* in protogalactic halo
- speculation: *differences in stellar mass function*?
- speculation:
  - low-mass halos → low level of turbulence → relatively massive stars
  - high-mass halos (atomic cooling halos) → high degree of turbulence → wider mass spectrum with peak at lower-masses?



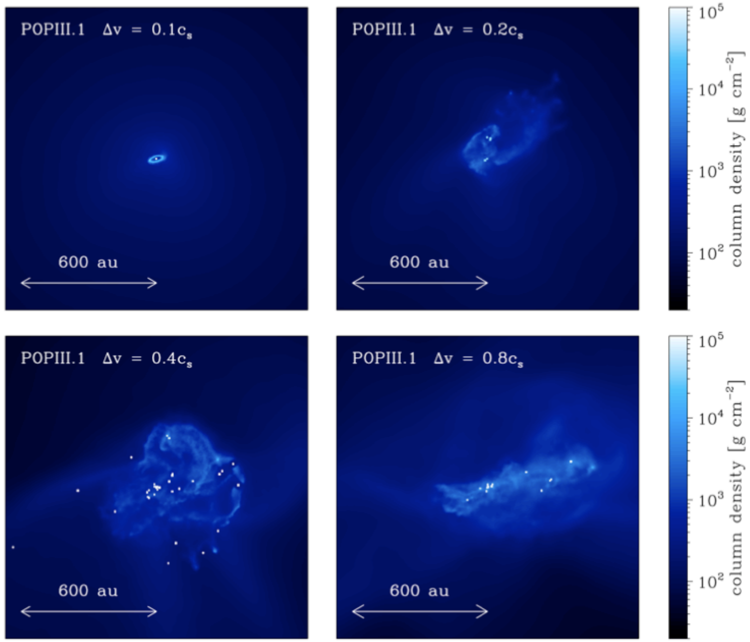
(Greif et al. 2008)

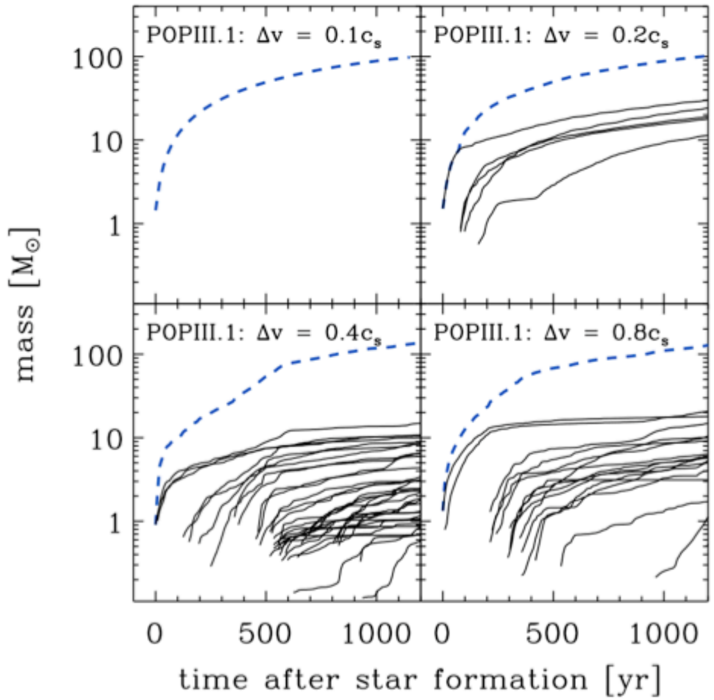


turbulence developing in an atomic cooling halo

(Greif et al. 2008)

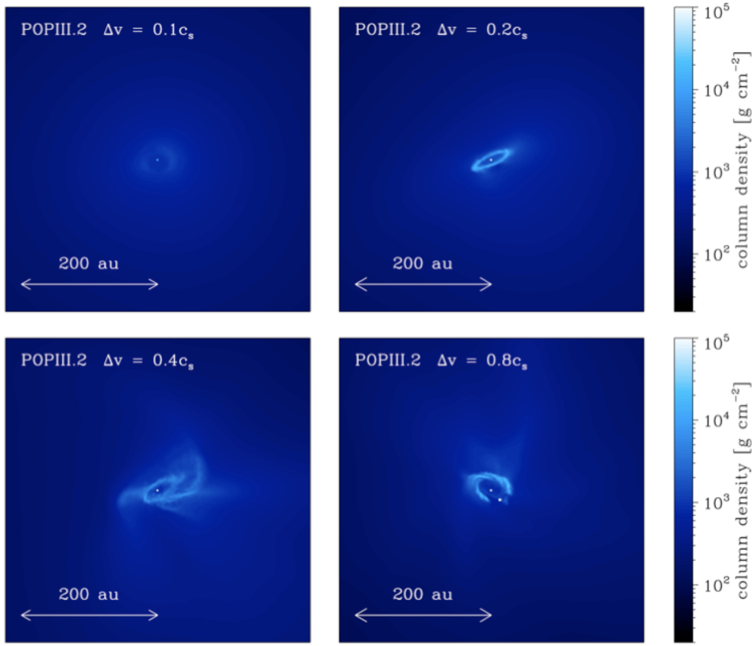
Pop III.1

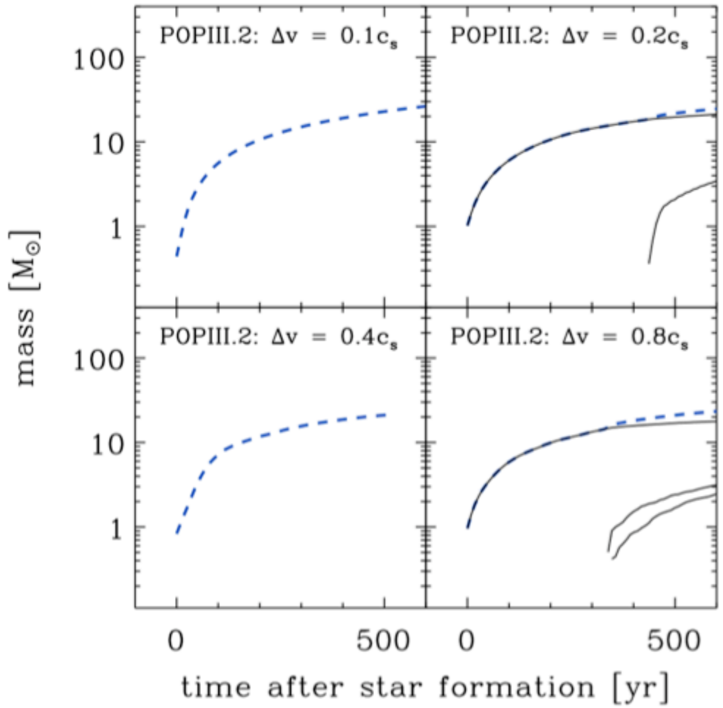




(Clark et al, submitted)

Pop III.2





(Clark et al, submitted)

once again: thermodynamics

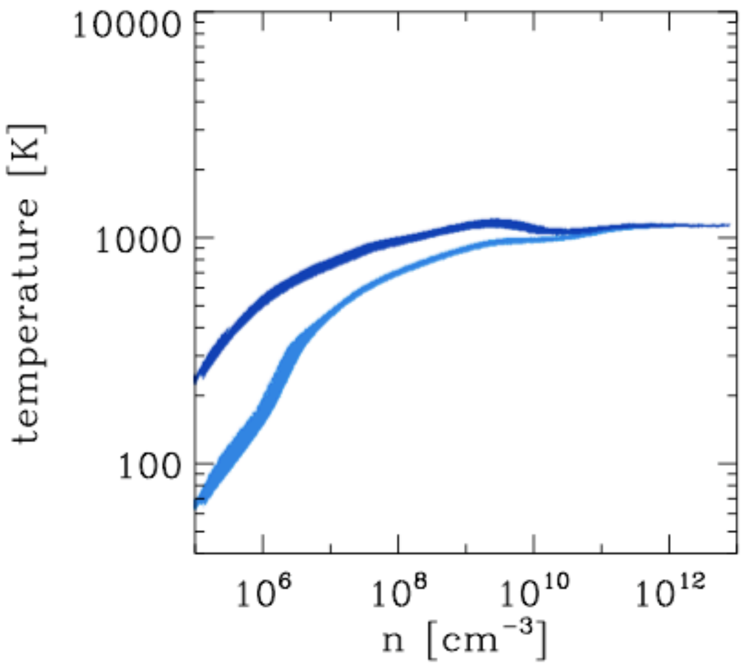


FIG. 6.— Temperature as a function of number density for the Pop. III.1 (dark blue) and Pop. III.2 (light blue) $\Delta v_{\rm turb} = 0.1\, c_{\rm s}$ simulations. In both cases, the curves denote the state of the cloud at the point just before the formation of the sink particle.

also Pop III.2 gas heats up above the CMB

--> weaker fragmentation!

once again: thermodynamics

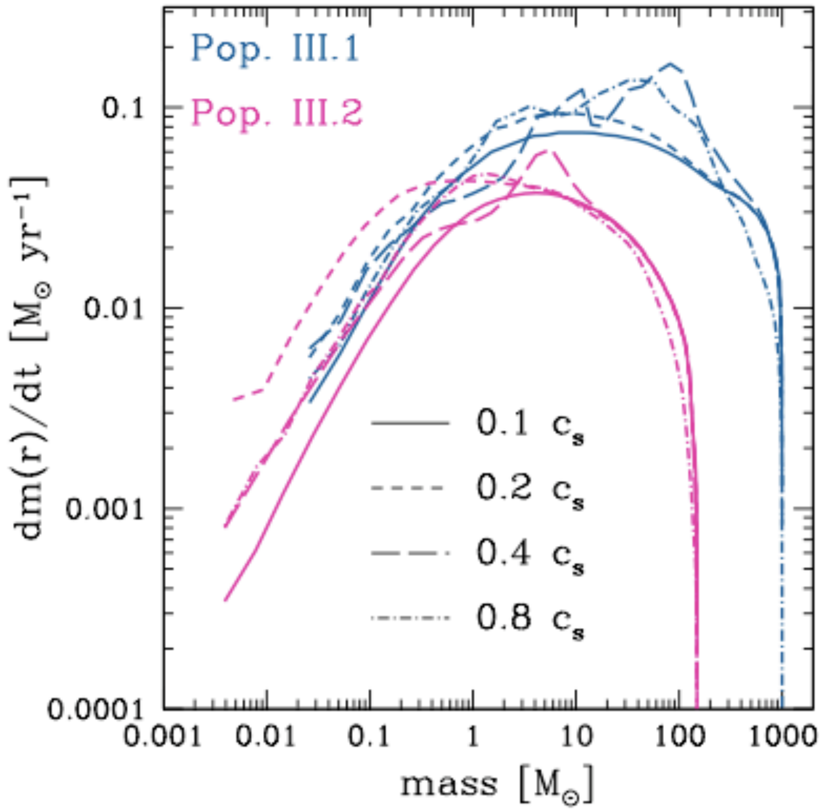
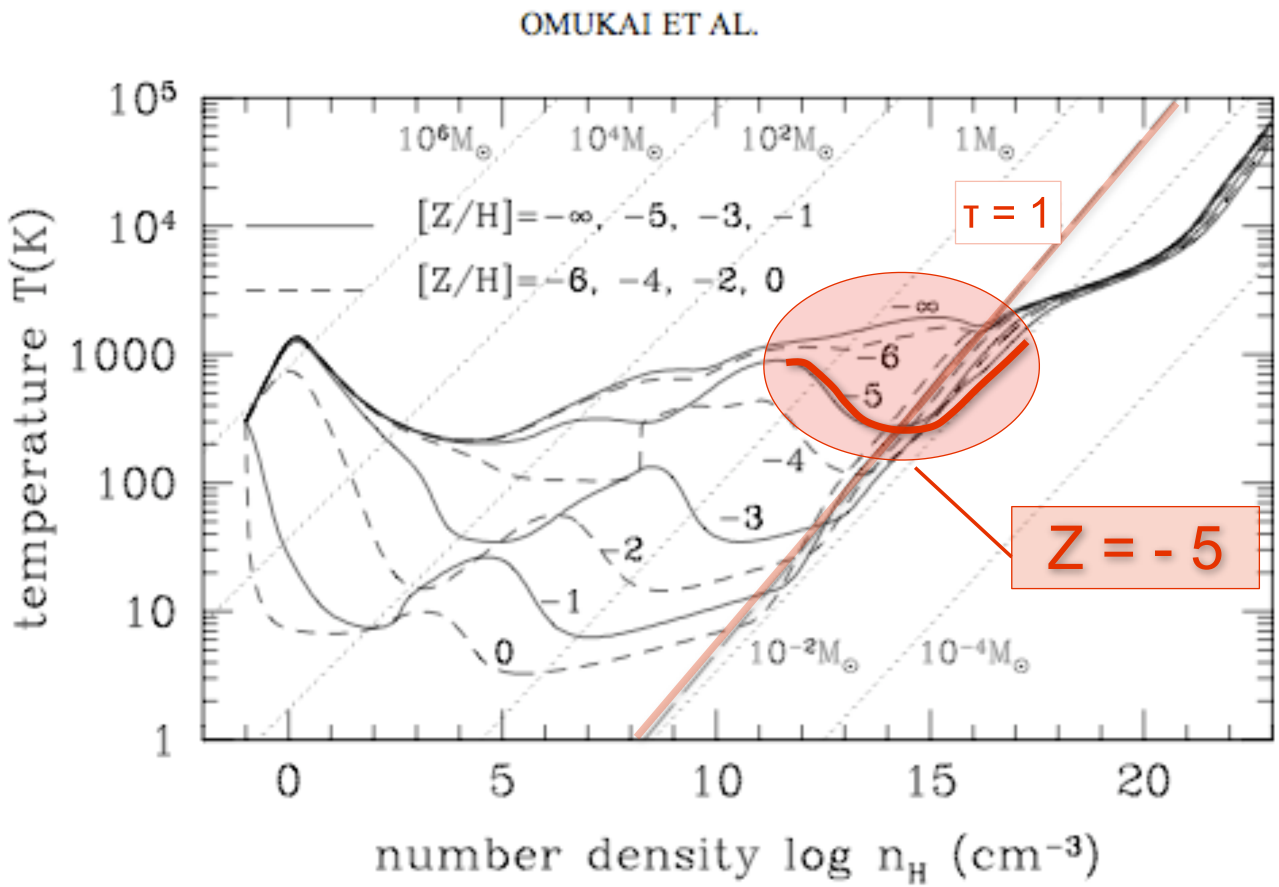


FIG. 8.— Accretion rates as a function of enclosed gas mass in the Pop. III.1 (upper lines; blue) and Pop. III.2 (lower lines; magenta) simulations, estimated as described in Section 4.1. Note that the sharp decline in the accretion rates for enclosed masses close to the initial cloud mass is an artifact of our problem setup; we would not expect to see this in a realistic Pop. III halo.

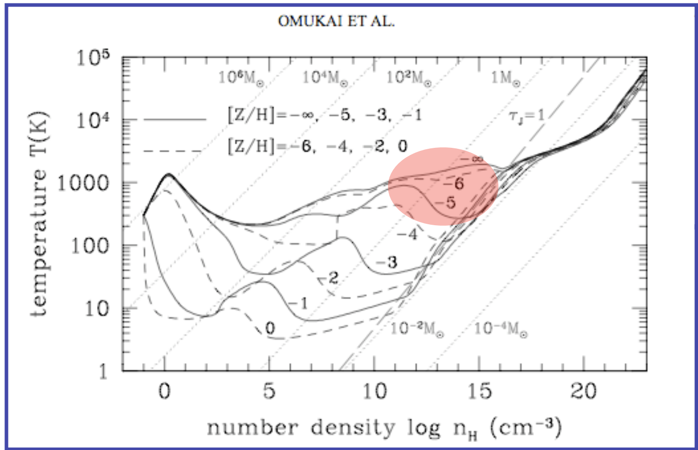
comparison of accretion rates...

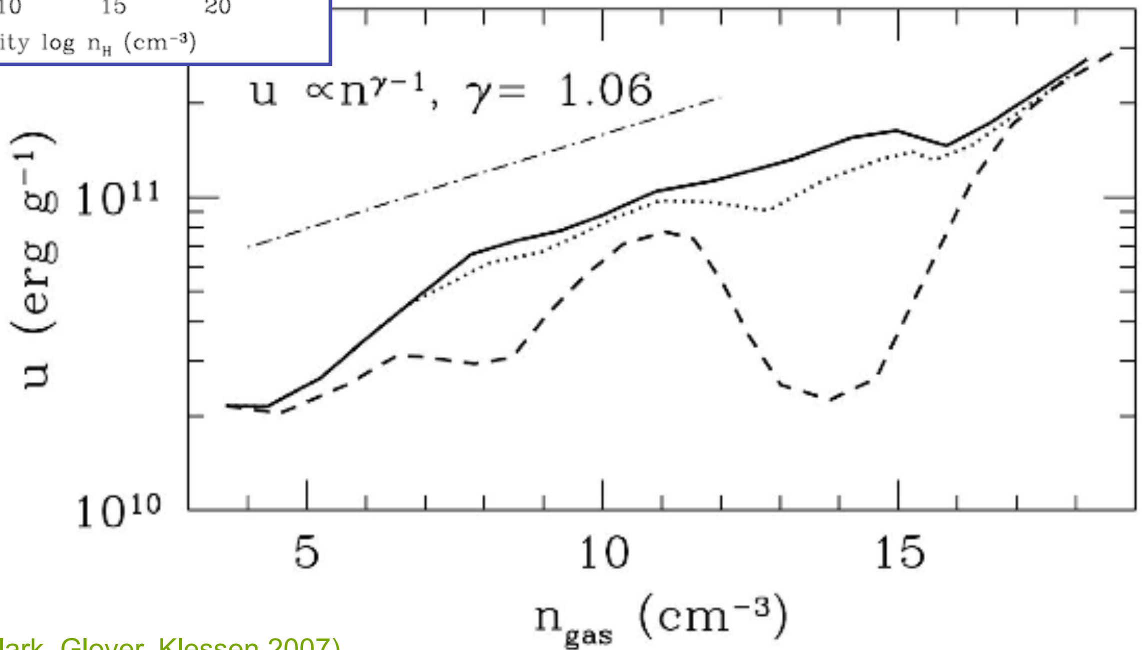
transition: Pop III to Pop II.5



(Omukai et al. 2005)

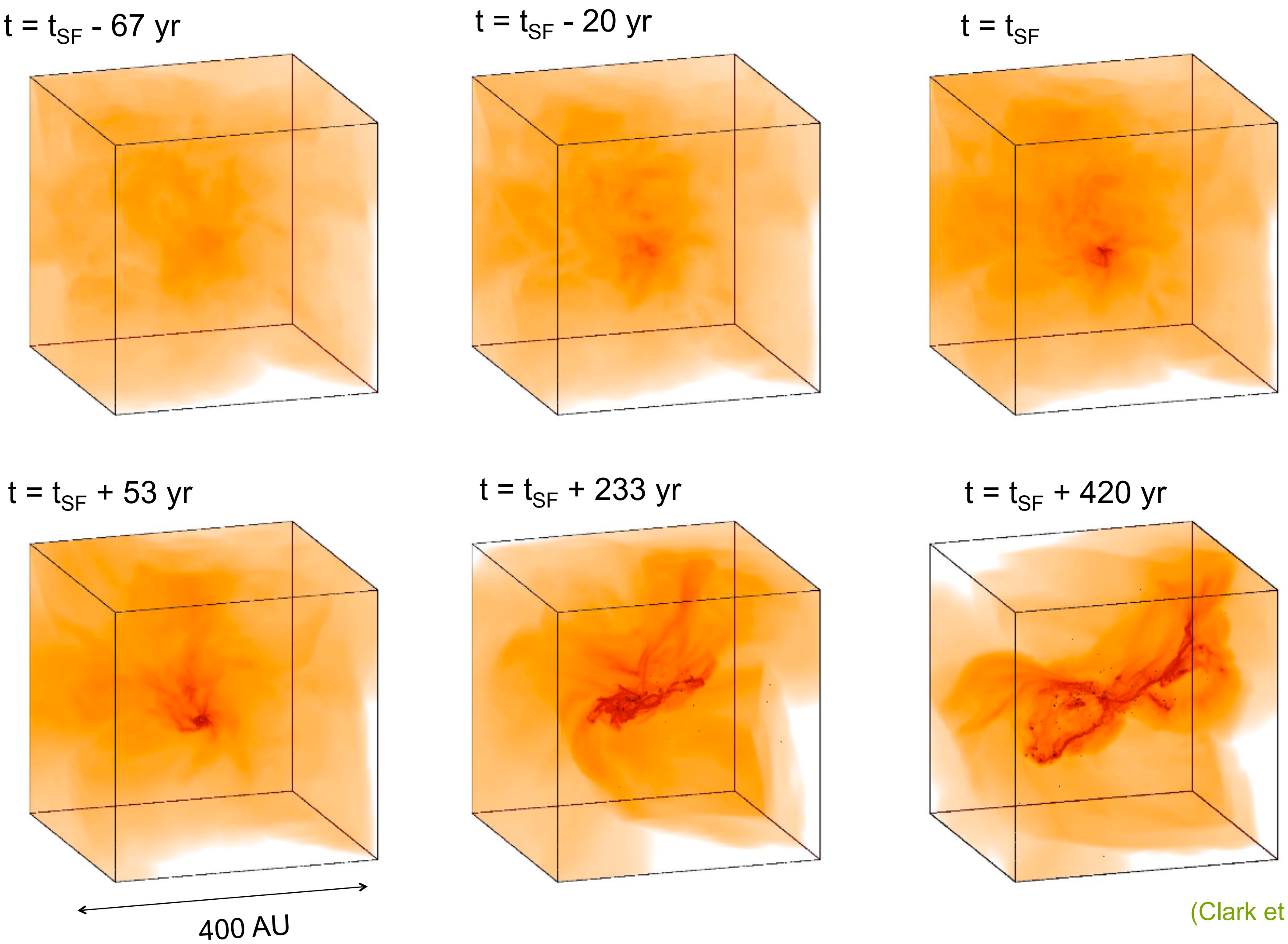
transition: Pop III to Pop II.5





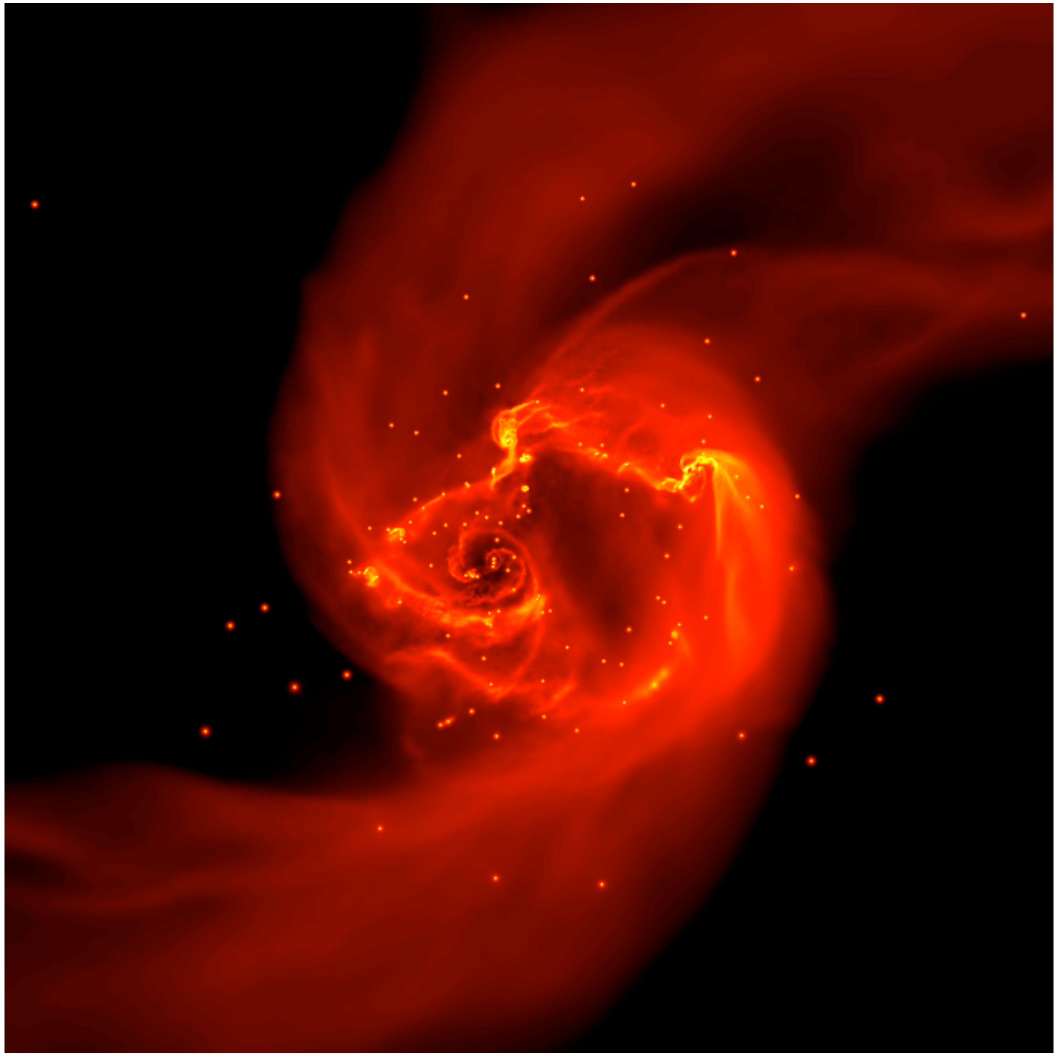
(Omukai et al. 2005, Clark, Glover, Klessen 2007)

dust induced fragmentation at $Z=10^{-5}$



(Clark et al. 2007)

dust induced fragmentation at $Z=10^{-5}$



dense cluster of low-mass protostars builds up:

- mass spectrum peaks *below 1 M_{sun}*
- cluster VERY dense $n_{stars} = 2.5 \times 10^9 \, pc^{-3}$
- fragmentation at density $n_{gas} = 10^{12} - 10^{13} \, cm^{-3}$

(Clark et al. 2008, ApJ 672, 757)

cluster build-up

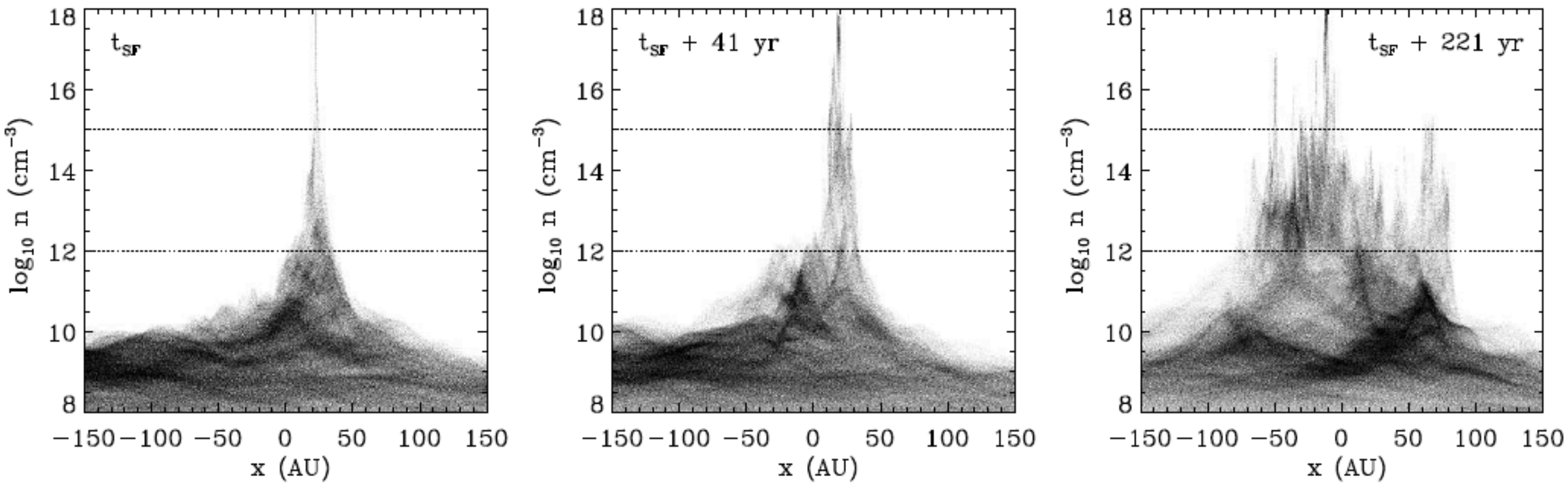
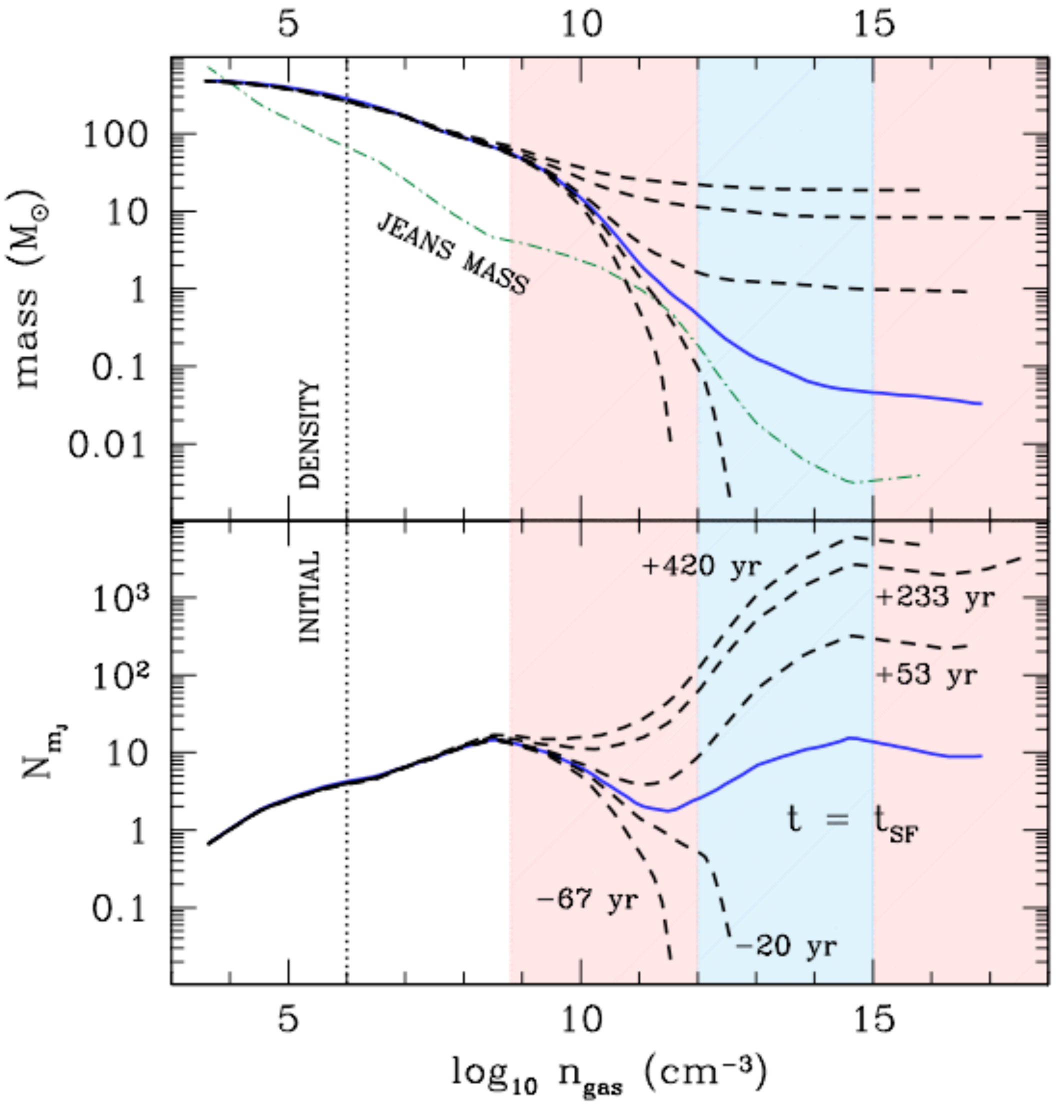


FIG. 3.— We illustrate the onset of the fragmentation process in the high resolution $Z = 10^{-5} Z_{\odot}$ simulation. The graphs show the densities of the particles, plotted as a function of their x-position. Note that for each plot, the particle data has been centered on the region of interest. We show here results at three different output times, ranging from the time that the first star forms (t_{sf}) to 221 years afterwards. The densities lying between the two horizontal dashed lines denote the range over which dust cooling lowers the gas temperature.

(Clark et al. 2007)

cluster build-up



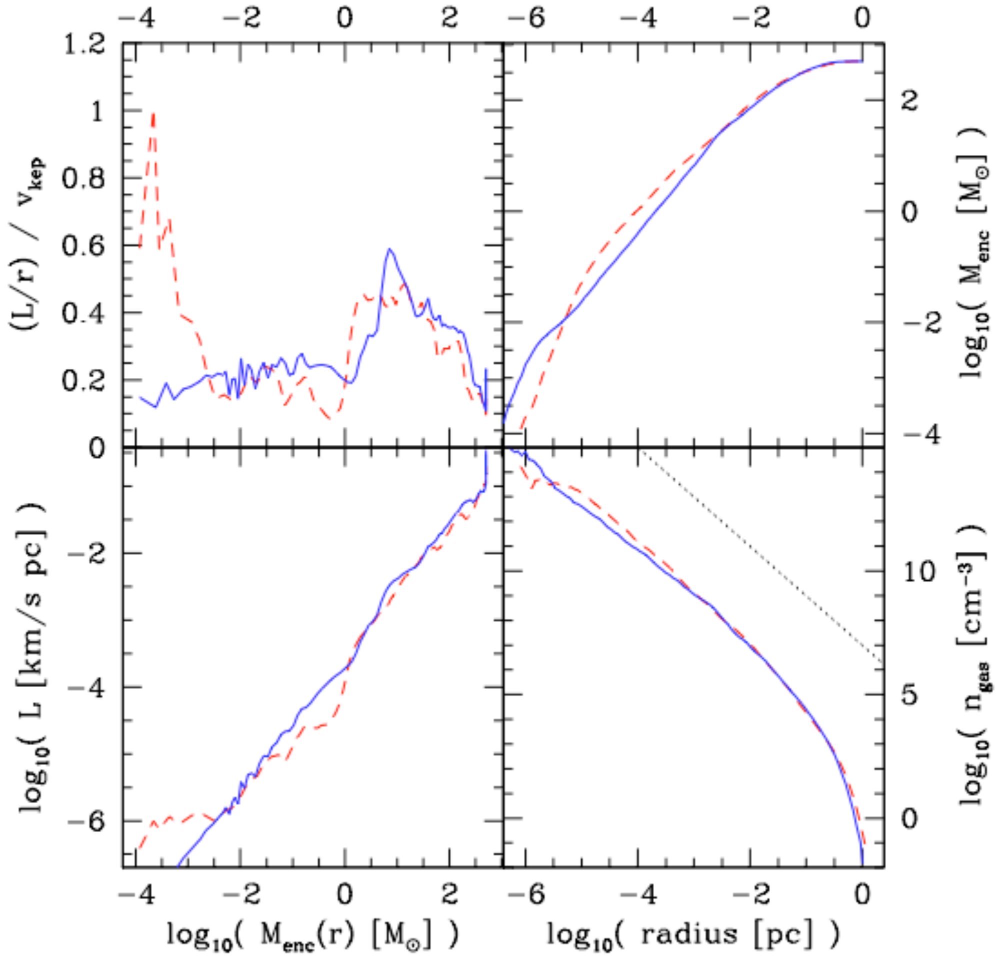
$\gamma > 1$
(heating)

$\gamma < 1$
(cooling)

(Clark et al. 2007)

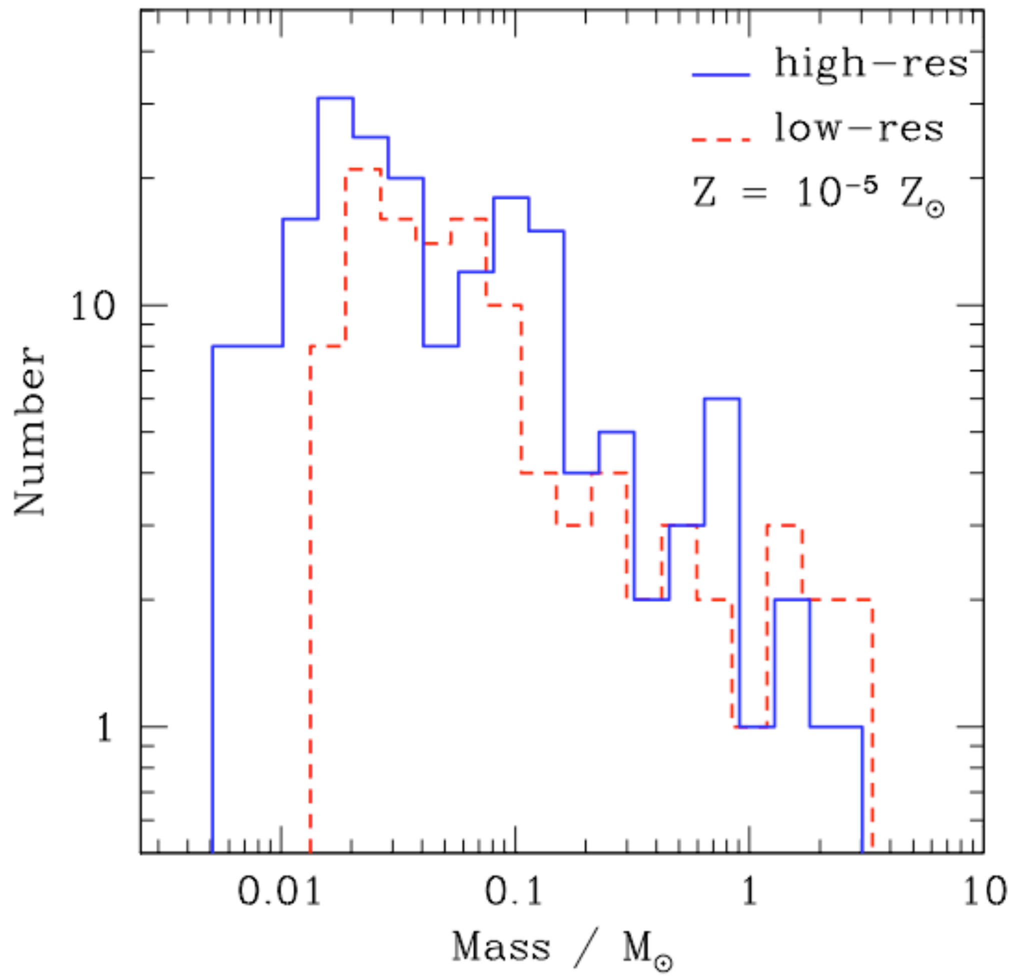
gas properties

gas properties at time when first star forms



(Clark et al. 2007)

dust induced fragmentation at $Z=10^{-5}$



dense cluster of low-mass protostars builds up:

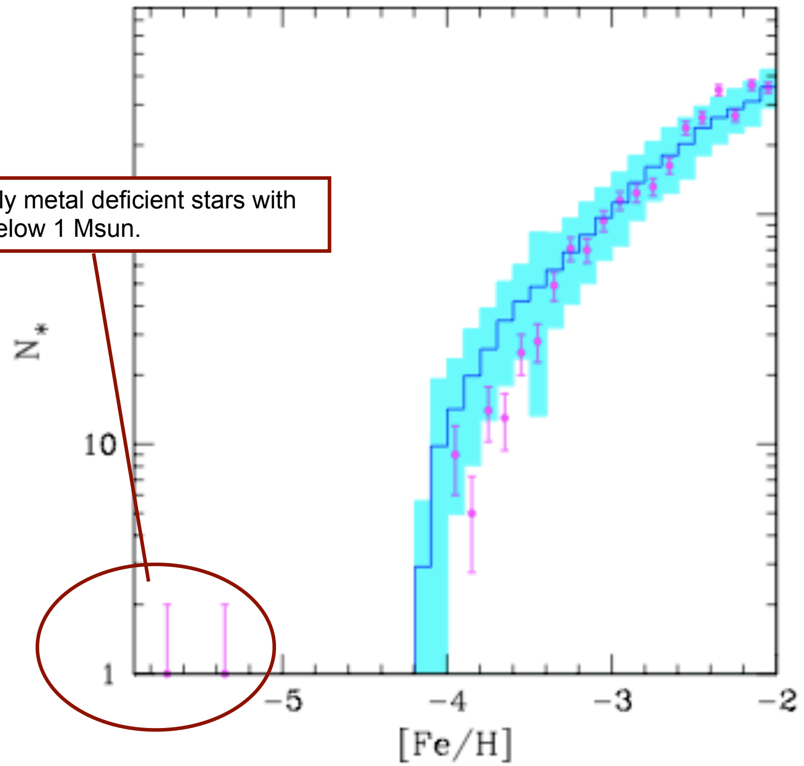
- mass spectrum peaks below 1 M_{sun}
- cluster VERY dense $n_{stars} = 2.5 \times 10^9 \, pc^{-3}$

- *predictions:*
  * low-mass stars with [Fe/H] ~ 10^{-5}
  * high binary fraction

(Clark et al. 2008)

dust induced fragmentation at Z=10^{-5}

2 extremely metal deficient stars with masses below 1 Msun.



(plot from Salvadori et al. 2006, data from Frebel et al. 2005)

dense cluster of low-mass protostars builds up:

- mass spectrum peaks below 1 M_{sun}
- cluster VERY dense n_{stars} = 2.5 x 10^9 pc^{-3}
- *predictions:*
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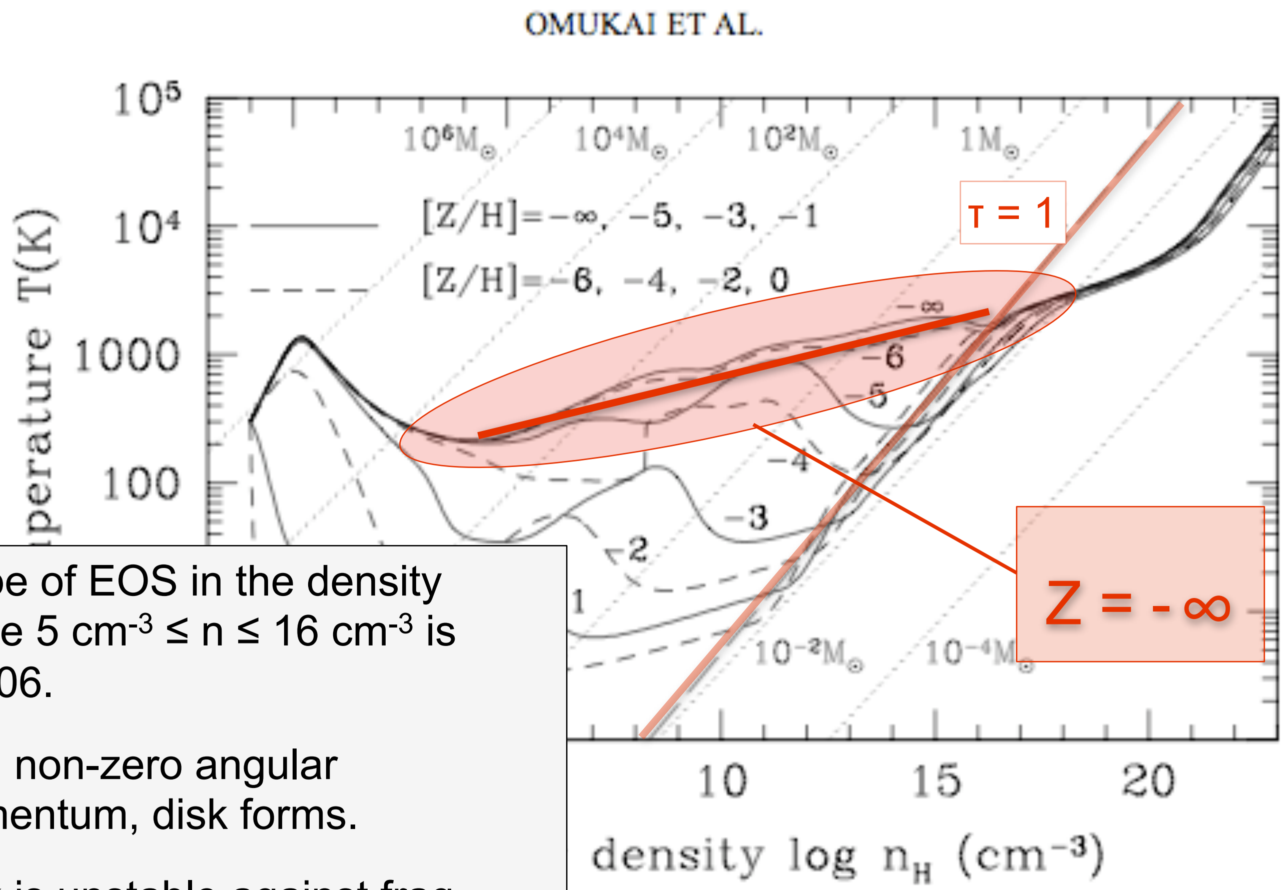
(Clark et al. 2008)

metal-free star formation



- slope of EOS in the density range 5 cm^{-3} ≤ n ≤ 16 cm^{-3} is γ≈1.06.
- with non-zero angular momentum, disk forms.
- disk is unstable against fragmentation at high density

more on Z=0 star formation

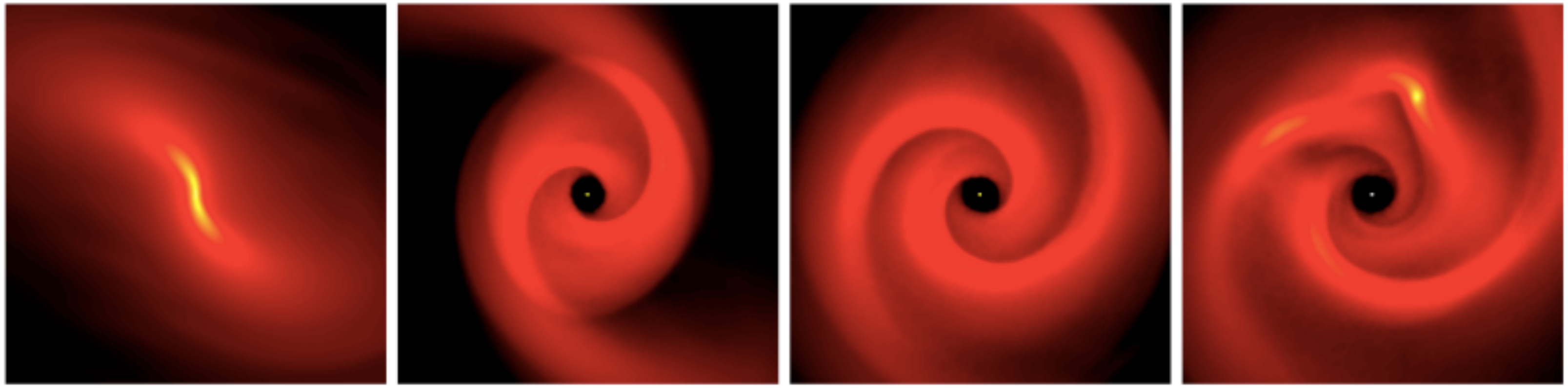


FIGURE 1. Column density images of the inner 66 au of the simulation, following the formation of the first protostar (sink particle) and the subsequent build-up of the protostellar disc and its eventual fragmentation. Starting from left-hand panel, which shows the gas at 1 yr before the protostar forms (t_{SF}), the next 3 panels show the evolution at times $t_{SF} + 76$ yr, $t_{SF} + 152$ yr and $t_{SF} + 228$ yr. The colour table is stretched from 10^3 g cm^{-2} to 10^6 g cm^{-2} .

more on Z=0 star formation

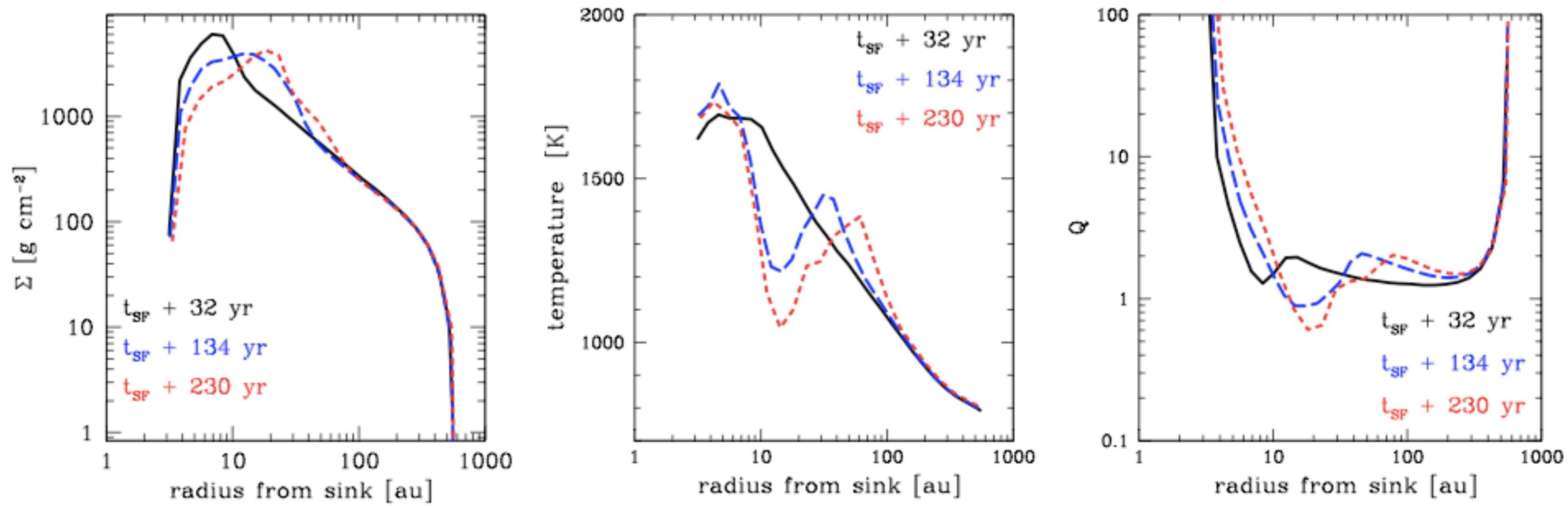


FIGURE 2. In the left-hand and central plots we show the radial profiles of the disc's surface density and gas temperature, centred on the first protostellar core to form in the simulation. The quantities are mass-weighted and taken from a slice through the midplane of the disc. In the right-hand plot we show the radial distribution of the corresponding Toomre parameter, $Q = c_s \kappa / \pi G \Sigma$, where c_s is the sound speed and κ is the epicyclic frequency. We adopt the standard simplification, and replace κ with the orbital frequency.

more on Z=0 star formation

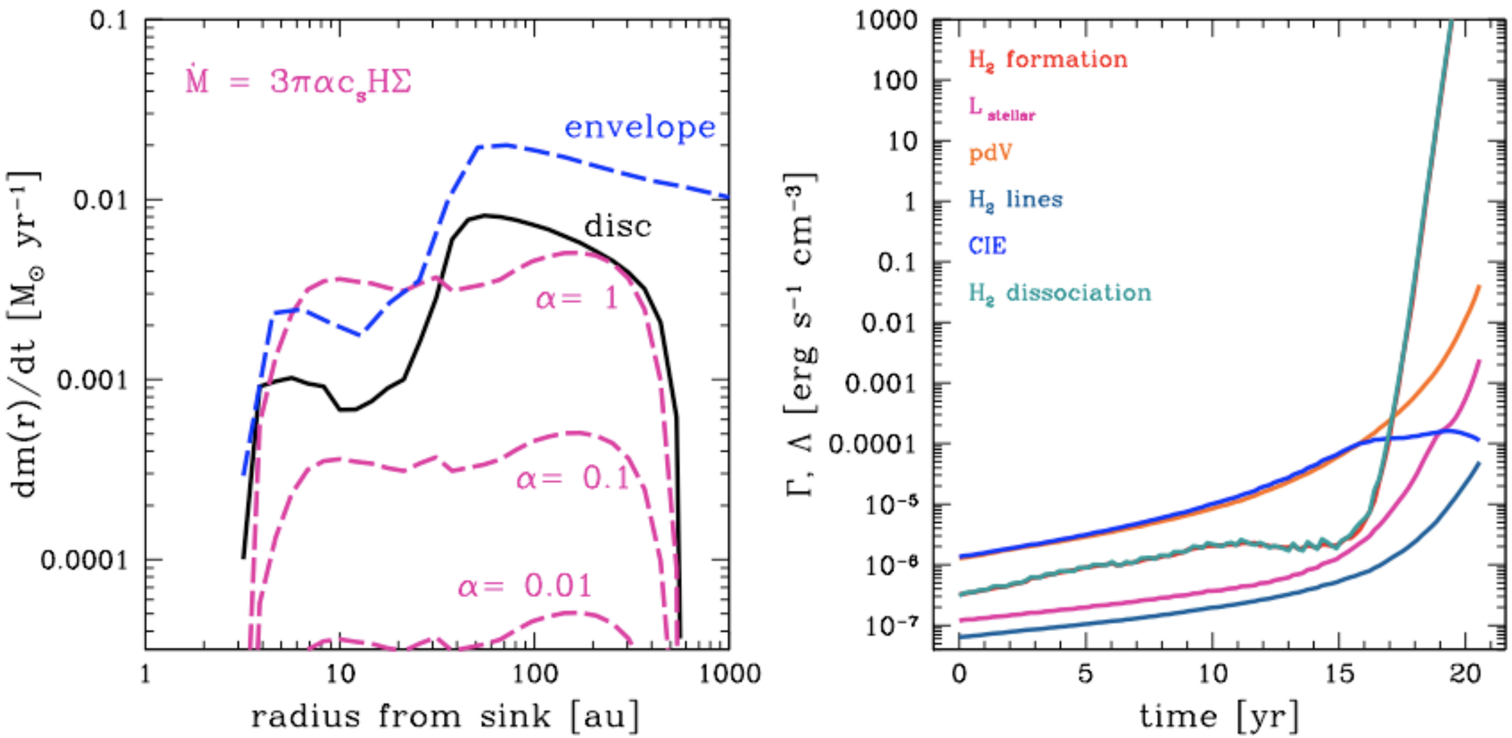
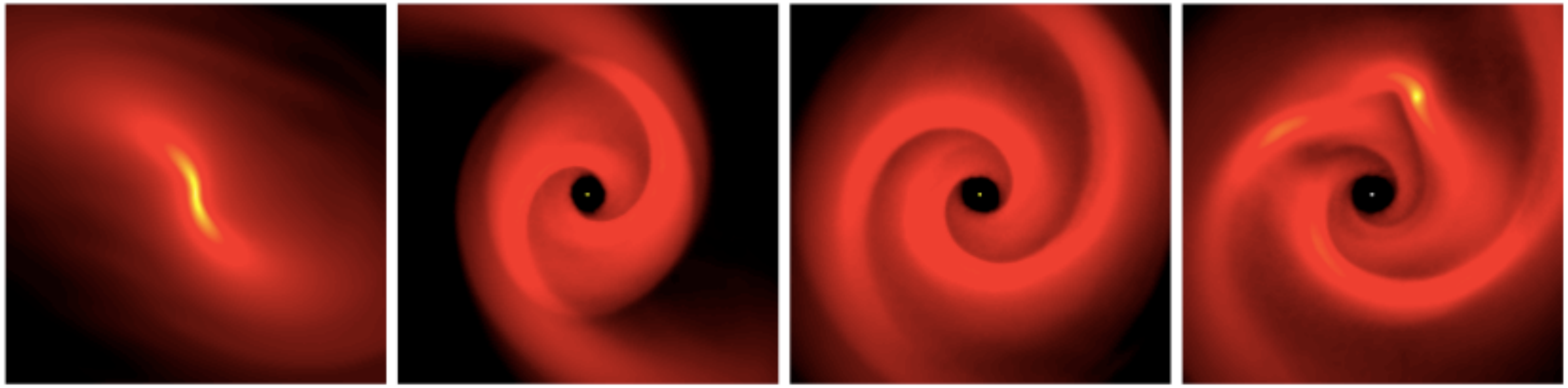


FIGURE 3. The left-hand plot shows the mass transfer through the disc. The solid black line shows the amount of mass moving inwards through each radial annulus in the disc per unit time. The dashed blue line shows the same quantity for the full spherical infalling envelope. The pink dashed lines show the accretion rates expected from an 'alpha' (thin) disc model, with three values of alpha. The right-hand plot shows the main heating and cooling processes that control the temperature evolution in the collapsing clump in the run-up to its eventual collapse.

primordial star formation



- first star formation is not less complex than present-day star formation
- brave claim: *all* Pop III stars form in multiple systems
- even braver claim: some Pop III stars fall in the mass range < 0.5 $M_\odot$ ---> they should still *be around*!!!!

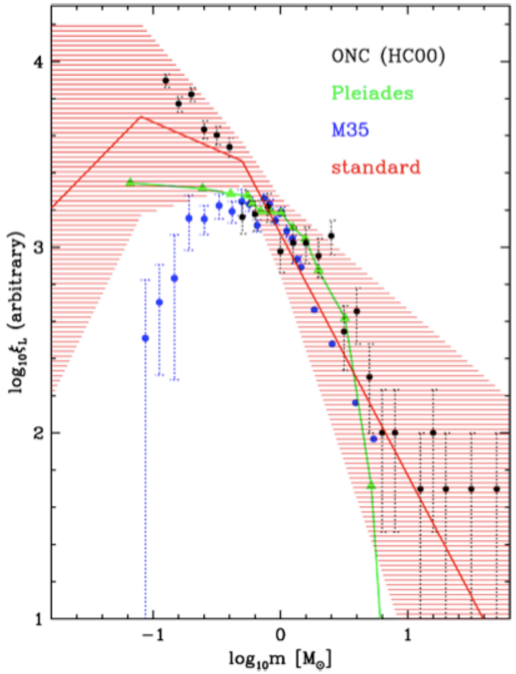
IMF

- distribution of stellar masses depends on
  - turbulent initial conditions
    --> mass spectrum of prestellar cloud cores
  - collapse and interaction of prestellar cores
    --> competitive accretion and *N*-body effects
  - *thermodynamic properties of gas*
    --> balance between heating and cooling
    --> EOS (determines which cores go into collapse)
  - (proto) stellar feedback terminates star formation
    ionizing radiation, bipolar outflows, winds, SN

Introduction

We want to address the following questions:

- What determines the upper stellar mass limit?
- What is the physics behind the observed HII regions?



IMF (Kroupa 2002)

Rosetta nebula (NGC 2237)

Feedback Processes

- radiation pressure on dust particles
- ionizing radiation
- stellar wind
- jets and outflows

Feedback Processes

- radiation pressure on dust particles
- ionizing radiation
- stellar wind
- jets and outflows

Radiation Pressure

has gained the most attention in the literature, most recent simulations by Krumholz et al. 2009

Ionization

only a few numerical studies so far (eg. Dale et al. 2007, Gritschneder et al. 2009), but H II regions around massive protostars can be observed!
$\longrightarrow$ direct comparison with observations possible

high-mass star formation

- focus on collapse of individual high-mass cores...
  - massive core with 1,000 $M_\odot$
  - Bonnor-Ebert type density profile
    (flat inner core with 0.5 pc and rho ~ $r^{-3/2}$ further out)
  - initial m=2 perturbation, rotation with β = 0.05
  - sink particle with radius 600 AU and threshold density of 7 x 10^{-16} g cm^{-3}
  - cell size 100 AU

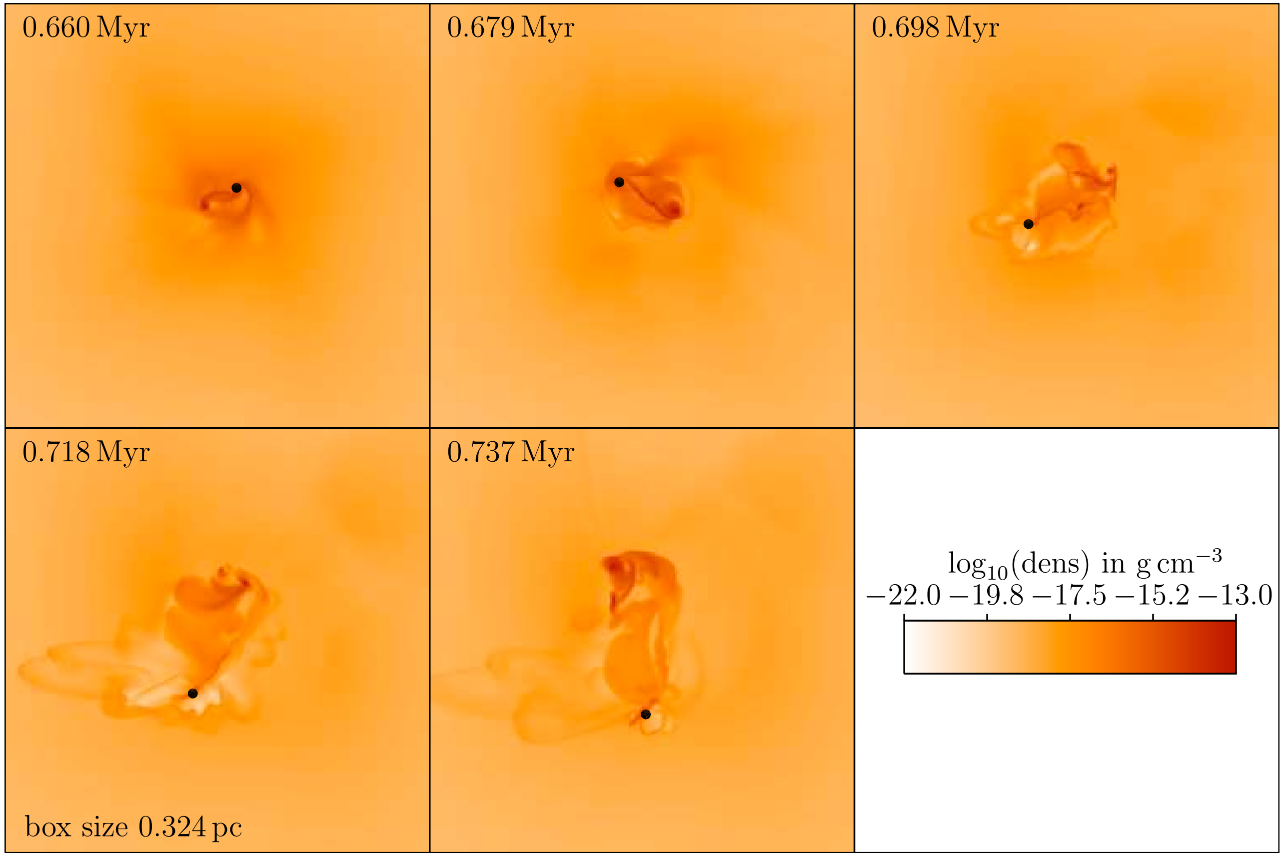
Peters et al. (2010a, ApJ, 711, 1017), Peters et al. (2010b, arXiv:1003.4998), Peters et al. (2010c,1005.3271)

high-mass star formation

- method:
  - FLASH with ionizing and non-ionizing radiation using raytracing based on hybrid-characteristics
  - protostellar model from Hosokawa & Omukai
  - rate equation for ionization fraction
  - relevant heating and cooling processes

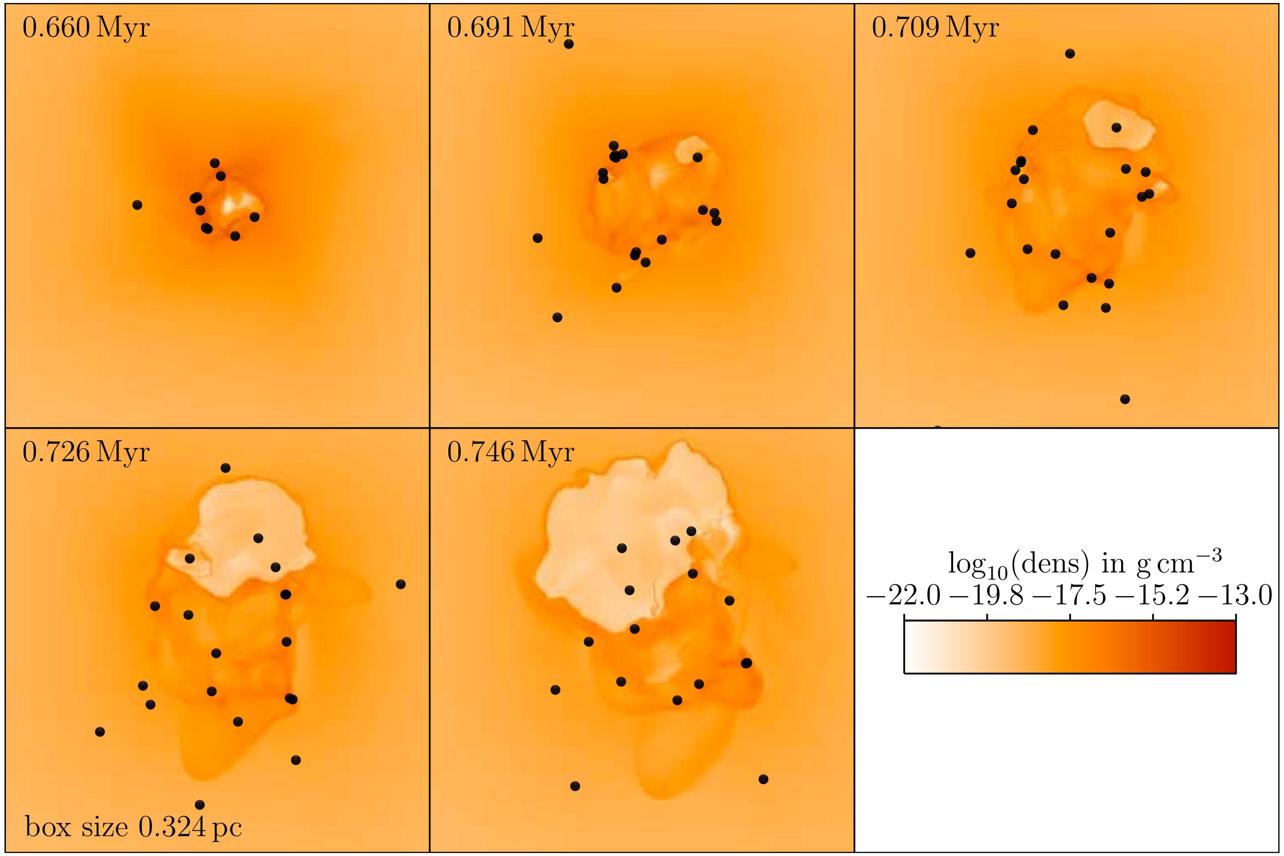
  - *first 3D calculations that consistently treat both ionizing and non-ionizing radiation in the context of high-mass star formation*

Peters et al. (2010a,b,c)



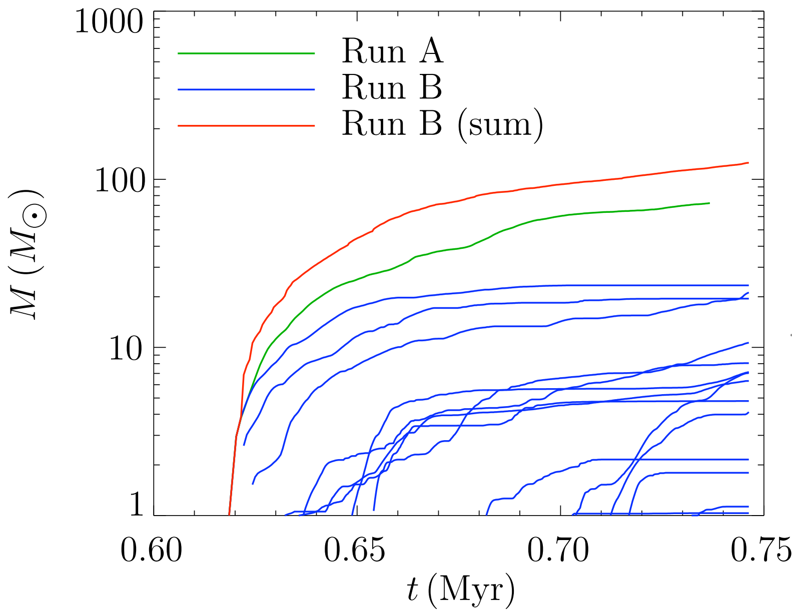
- disk is gravitationally unstable and fragments
- we suppress secondary sink formation by "Jeans heating"
- H II region is shielded effectively by dense filaments
- ionization feedback does not cut off accretion!

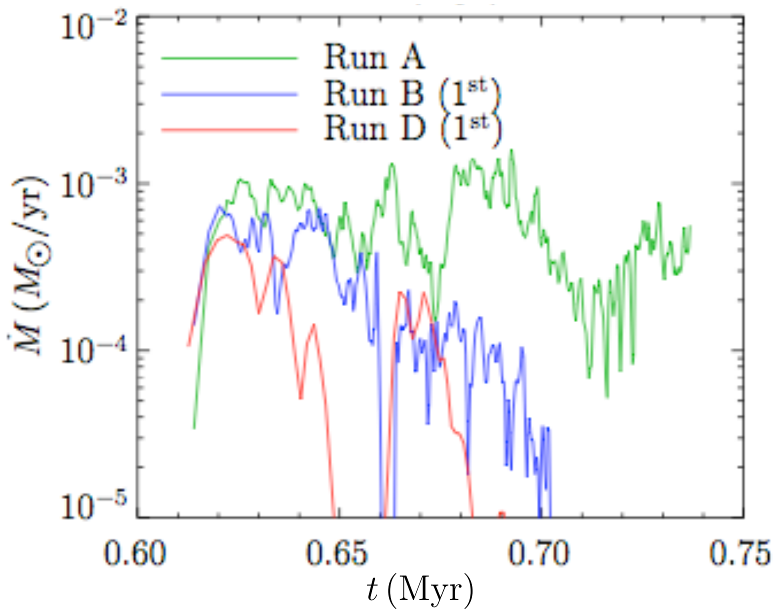
Peters et al. (2010a,b,c)



- all protostars accrete from common gas reservoir
- accretion flow suppresses expansion of ionized bubble
- cluster shows "fragmentation-induced starvation"
- halting of accretion flow allows bubble to expand

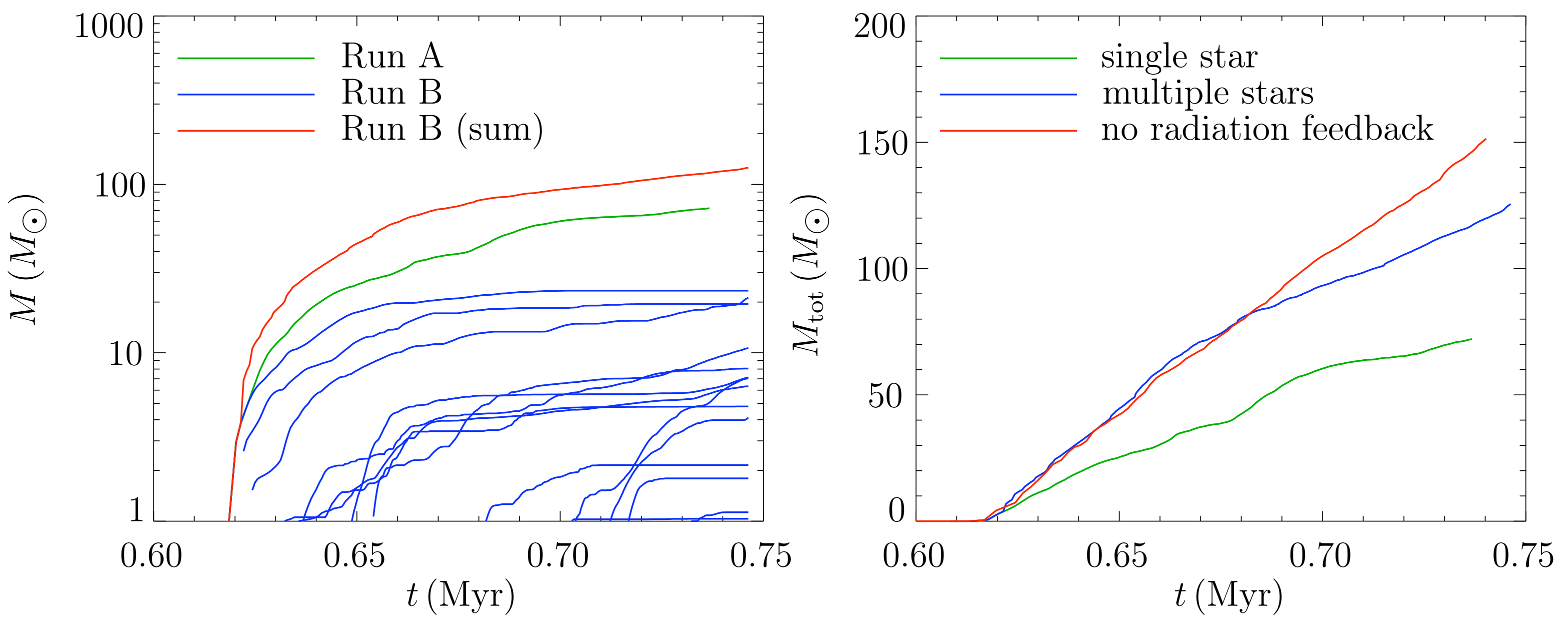
Peters et al. (2010a,b,c)





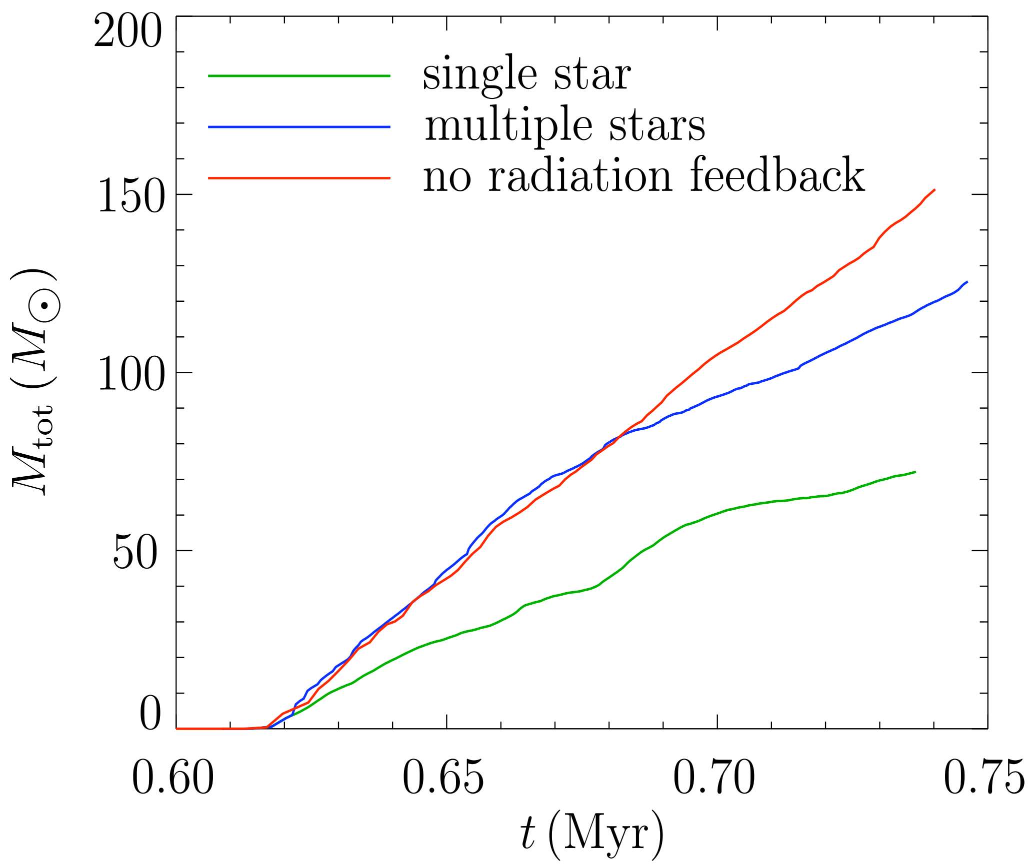
- single protostar accretes $72 M_\odot$ in $120\,\mathrm{kyr}$ (Run A)
- ionization feedback alone is unable to stop accretion
- accretion is limited when multiple protostars can form (Run B)
- no star in multi sink simulation reaches more than $30 M_\odot$

Peters et al. (2010a,b,c)



- compare with control run without radiation feedback
- total accretion rate does not change with accretion heating
- expansion of ionized bubble causes turn-off
- no triggered star formation by expanding bubble

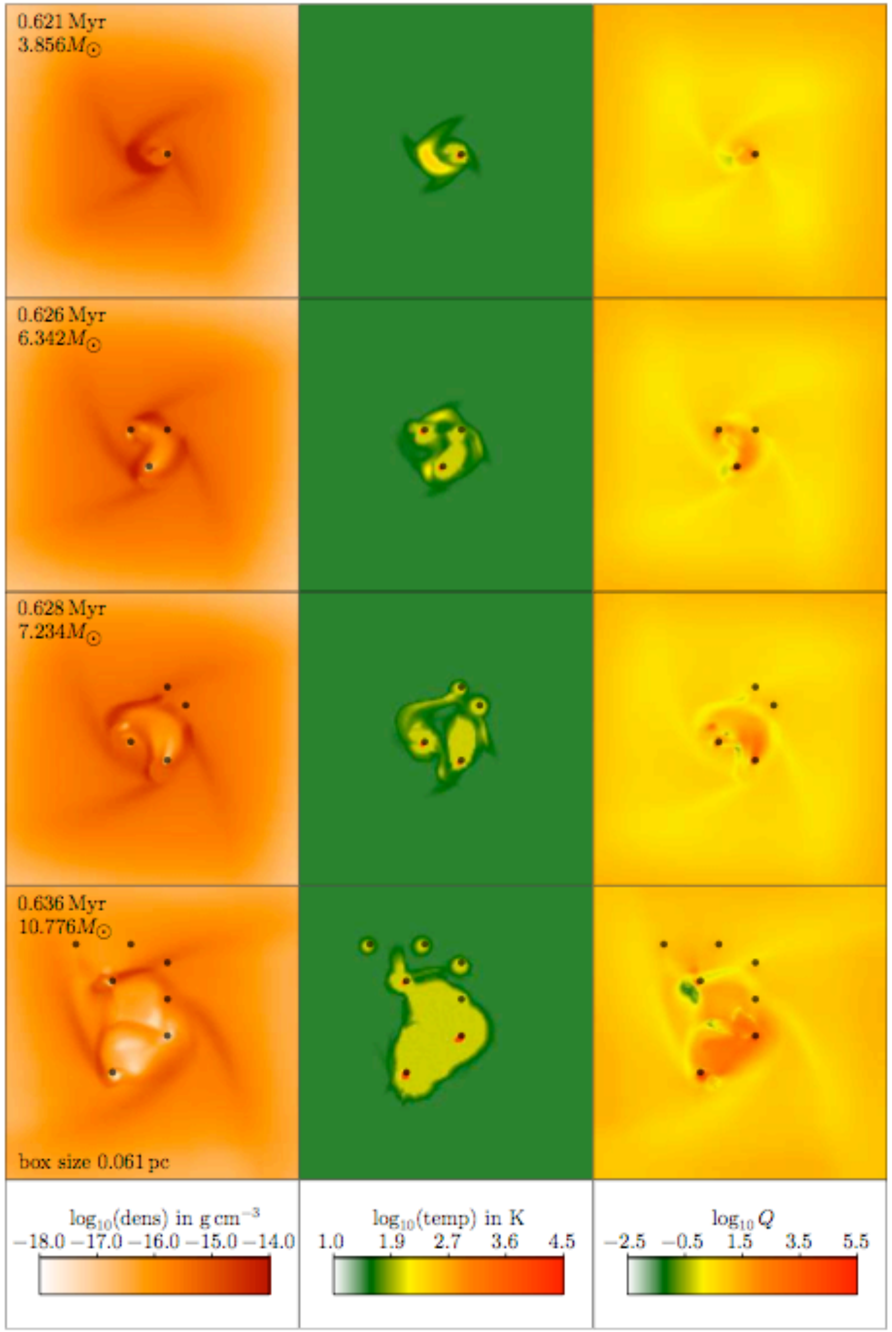
Peters et al. (2010a,b,c)

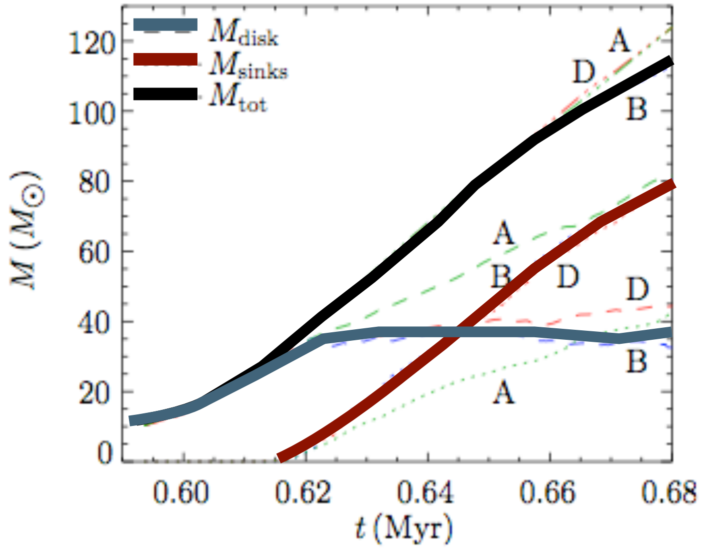


OVERVIEW OF COLLAPSE SIMULATIONS.

Name	Resolution	Radiative Feedback	Multiple Sinks	$M_{\rm sinks}$ ($M_\odot$)	$N_{\rm sinks}$	$M_{\rm max}$ ($M_\odot$)
Run A	98 AU	yes	no	72.13	1	72.13
Run B	98 AU	yes	yes	125.56	25	23.39
Run D	98 AU	no	yes	151.43	37	14.64

Peters et al. (2010a,b,c)

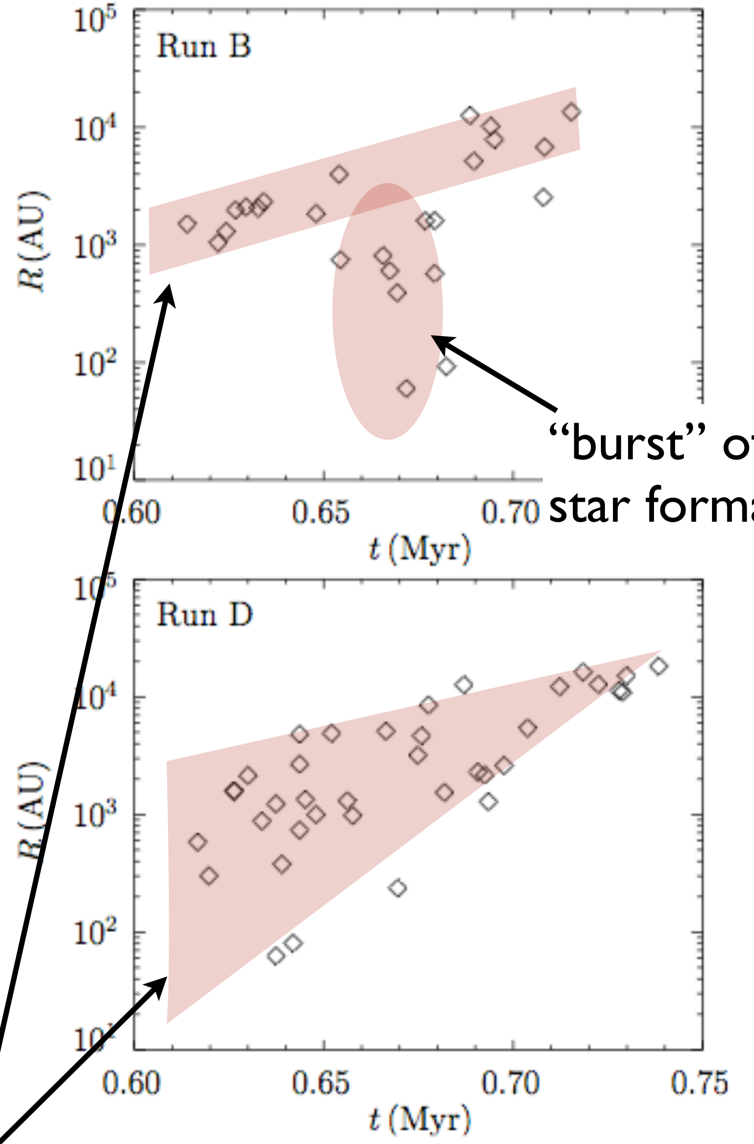




mass load onto the disk exceeds inward transport --> becomes gravitationally unstable (see also Kratter & Matzner 2006, Kratter et al. 2010)

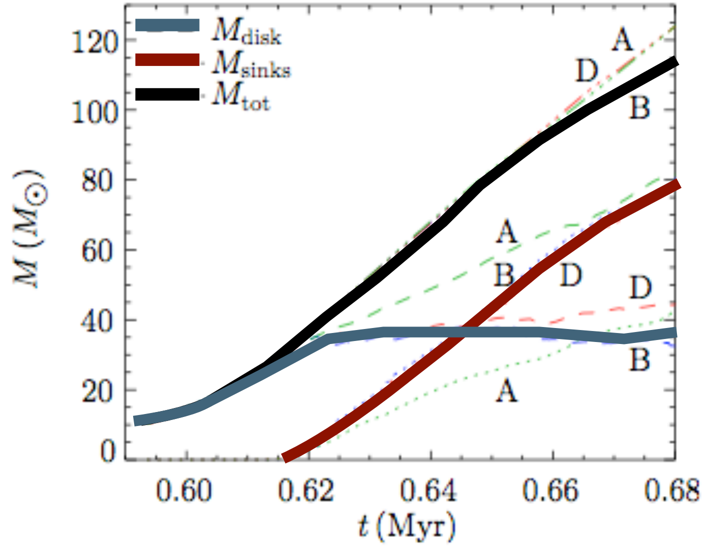
fragments to form multiple stars --> explains why high-mass stars are seen in clusters

Peters et al. (2010a,b,c)



"burst" of star formation

younger protostars form at larger radii



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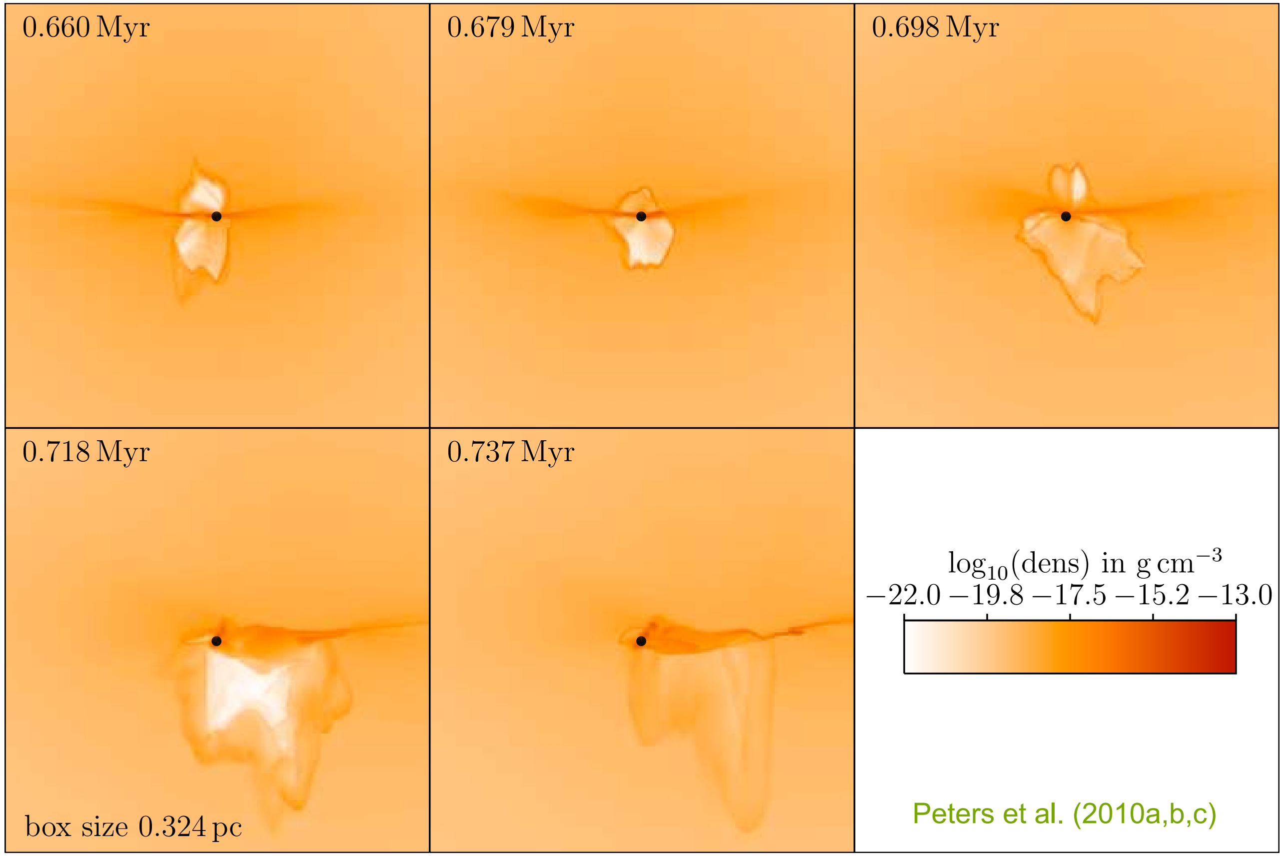
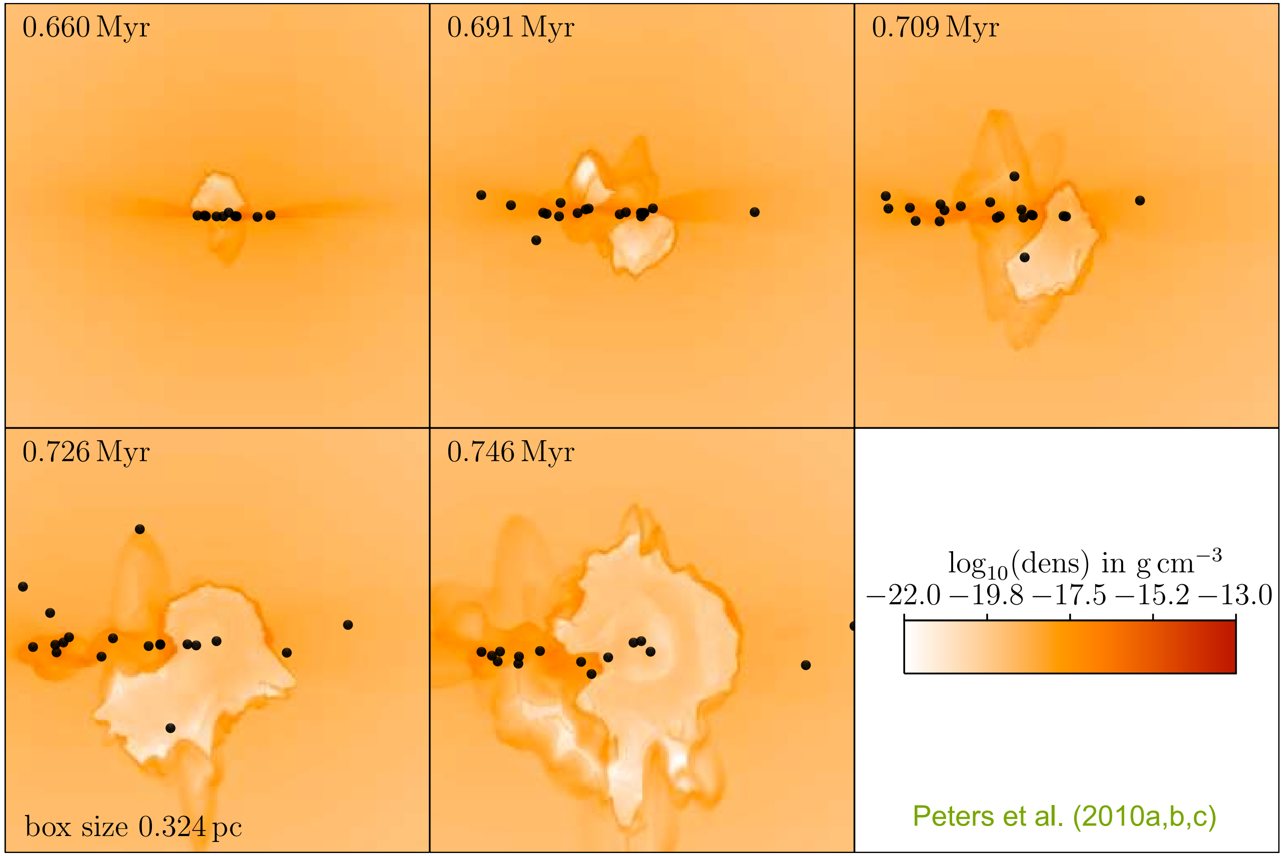


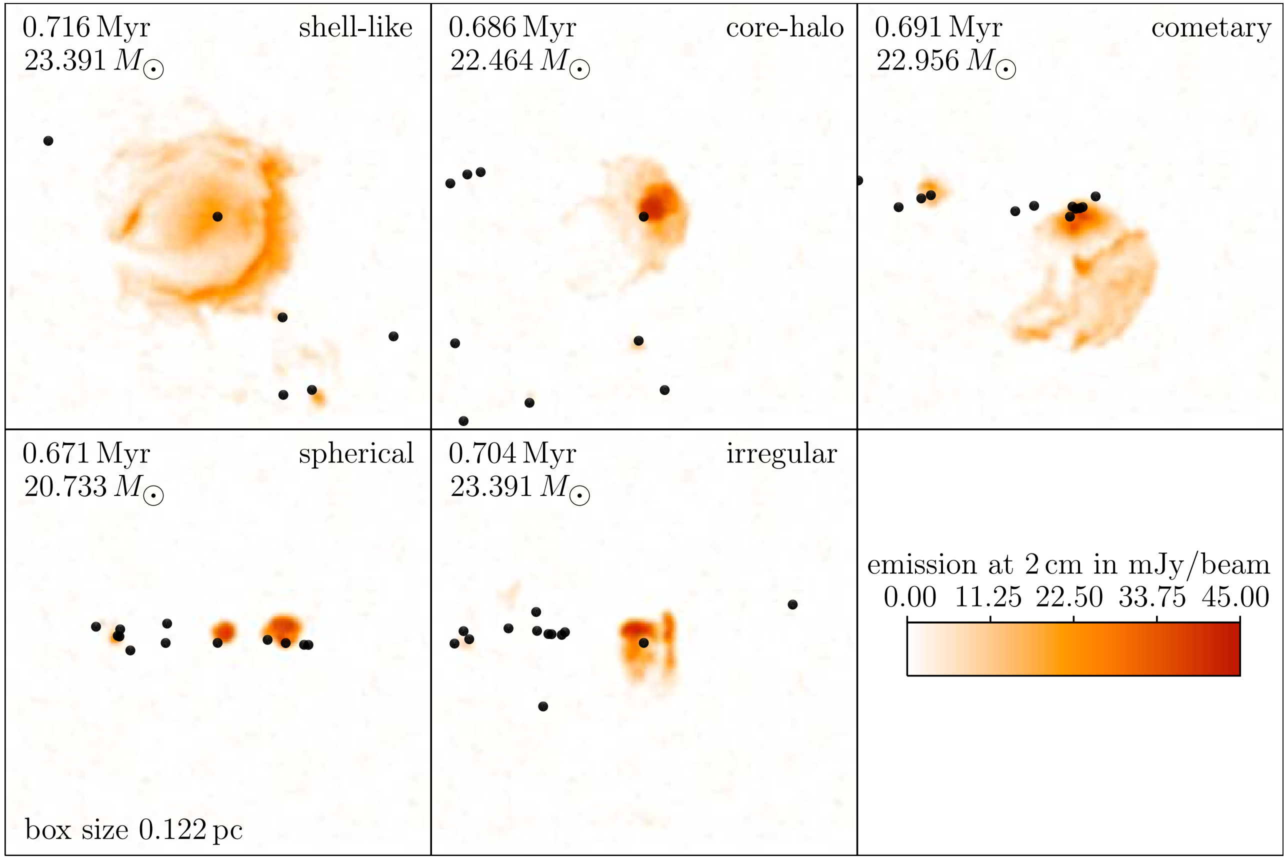
- thermal pressure drives bipolar outflow
- filaments can effectively shield ionizing radiation
- when thermal support gets lost, outflow gets quenched again
- no direct relation between mass of star and size of outflow



- bipolar outflow during accretion phase
- when accretion flow stops, ionized bubble can expand
- expansion is highly anisotropic
- bubbles around most massive stars merge

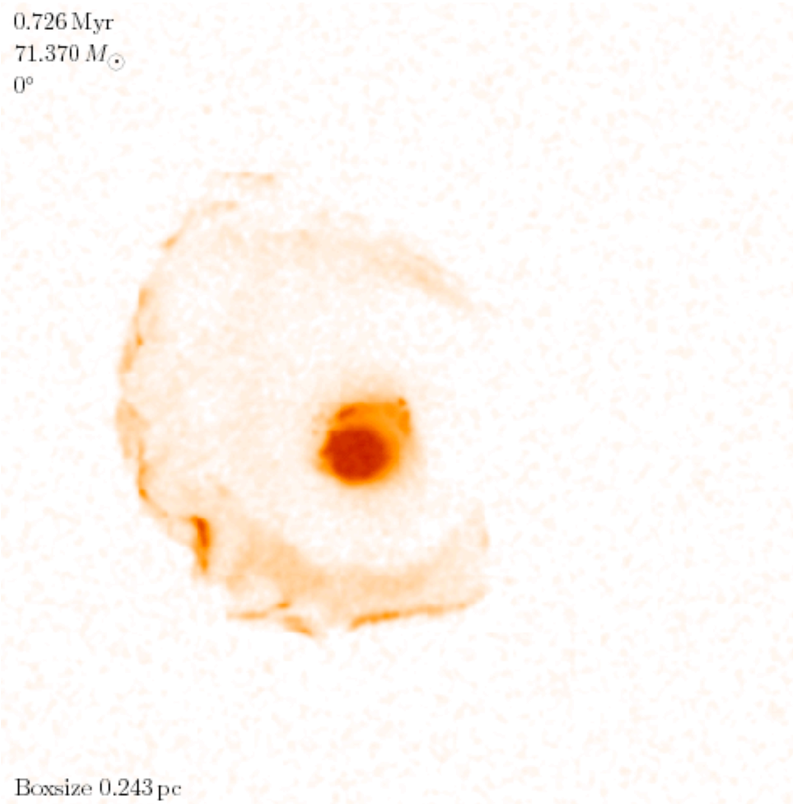
- numerical data can be used to generate continuum maps
- calculate free-free absorption coefficient for every cell
- integrate radiative transfer equation (neglecting scattering)
- convolve resulting image with beam width
- VLA parameters:
  - distance $2.65\,\mathrm{kpc}$
  - wavelength $2\,\mathrm{cm}$
  - FWHM $0\farcs14$
  - noise $10^{-3}\,\mathrm{Jy}$

Peters et al. (2010a,b,c)



- synthetic VLA observations at $2\,\mathrm{cm}$ of simulation data
- interaction of ionizing radiation with accretion flow creates high variability in time and shape
- flickering resolves the lifetime paradox!

Peters et al. (2010a,b,c)



Morphology of HII region depends on viewing angle

Peters et al. (2010a,b,c)

H II Region Morphologies

Type	WC89	K94	single	multiple
Spherical/Unresolved	43	55	19	60 ± 5
Cometary	20	16	7	10 ± 5
Core-halo	16	9	15	4 ± 2
Shell-like	4	1	3	5 ± 1
Irregular	17	19	57	21 ± 5

WC89: Wood & Churchwell 1989, K94: Kurtz et al. 1994

- statistics over 25 simulation snapshots and 20 viewing angles
- statistics can be used to distinguish between different models
- single sink simulation does not reproduce lifetime problem

Peters et al. (2010a,b,c)

Conclusions and Outlook

Conclusions

- Ionization feedback cannot stop accretion
- Ionization drives bipolar outflow
- H II region shows high variability in time and shape
- All classified morphologies can be observed in one run
- Lifetime of H II region determined by accretion time scale
- Rapid accretion through dense, unstable flows
- Fragmentation-induced mass limits of massive stars

star formation

- stars form in clusters (at all cosmic ages!)
- star formation is a highly complex process, involving multiple scales and multiple physical processes
- initial conditions matter big time
- first star formation is not less complex than present-day star formation
- IMF is result of many processes (turbulence, *N*-body dynamics, thermodynamics, feedback, etc.)
- IMF is "easy" to get, look for secondary statistics